On the origin of cyclical variability in the winds of massive stars

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On the origin of cyclical variability in the winds of massive stars

Jeroen A. de Jong
On the origin of cyclical variability in the winds of massive stars

Over de oorsprong van de cyclische veranderlijkheid in de winden van zware sterren

*Academisch Proefschrift*

ter verkrijging van de graad van doctor aan de Universiteit van Amsterdam op gezag van de Rector Magnificus prof. dr J.J.M. Franse ten overstaan van een door het college voor promoties ingestelde commissie in het openbaar te verdedigen in de Aula der Universiteit op dinsdag 8 februari 2000 te 15:00 uur

doors

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geboren te Beverwijk
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The research described in this thesis was supported by the Netherlands Foundation for Research in Astronomy (NFRA), with financial aid from the Netherlands Organisation for Scientific Research (NWO) under project 781-71-053.

ISBN 90-5776-043-6
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Introduction and summary

1.1. Introduction

The goal of this thesis work is the search for the origin of the widely observed cyclical variability of winds of hot, massive stars. This is one of the most challenging problems in the research of massive stars during the past 15 years. We first describe the considerations that lead us to this investigation, and highlight in the following sections the background and importance of this research.

At present, two different mechanisms are proposed as a possible cause of the variability of winds of massive stars: non-radial pulsations (where the phase of the oscillation varies over the surface) and magnetic fields, both of which being surface phenomena that are very difficult to observe in these stars. The observational support is very scarce. At present there are only 6 confirmed O stars with non-radial pulsations, two of which were found during this thesis work (as the result of 8 years of progress in theory and techniques of analysis after the data had been obtained), and there is no O star at all with a magnetic field detection. However, the evidence that the periodicity of the wind variability is directly related to the rotation period of the star is overwhelming. This implies that some surface phenomenon must play a dominant role in modulating the wind, and the presence of (small) magnetic structures anchored to the surface is the most obvious candidate, although non-radial pulsation modes can in principle do the same job.

Non-radial pulsations are expected to occur in basically all stars. The theory predicts this kind of instabilities over a large part of the HR diagram, but the predictions of the pulsation periods or amplitudes for hot, massive stars are only qualitative so far. It is difficult to find non-radial pulsations in these stars because non-radial pulsations can only be detected as moving, very small perturbations in the line profiles, often less than 0.1% of the intensity. This requires 2m-class telescopes equipped with high-dispersion spectrographs, even for the brightest stars. The observational problem is complicated even more because the timescales for pulsation, outflow and rotation are all of the order of hours to days, and any serious attempt to study these phenomena requires the simultaneous effort of several observatories around the globe. Presently the (Fourier) techniques to detect non-radial pulsations in rotating stars are reasonably well developed (see Telling 1996, Schrijvers 1999), but the bottleneck remains to obtain a homogeneous dataset with sufficient coverage (typically a few observations per hour) over a long enough time interval (typically a week).

Although at first sight it may seem obvious that magnetic fields must play a major role in modulating the outflowing stellar wind, the difficulties with magnetic fields are manyfold. First of all, there is no theory that predicts the occurrence of a specific magnetic field configuration on the surface of massive stars, mainly because these stars have no outer convection layer such as in solar-type stars. The observational problems are at least as severe. One needs a specialized instrument of which there are only very few available. The present detection techniques, based on the Zeeman effect, make use of the fact that the difference in wavelength of the oppositely shifted Zeeman components of a spectral line is a measure of the strength of the magnetic field. For small fields this difference is at the border of what can be detected above the noise. Modern multiplexing techniques use the accumulated signal of all the observed lines together. This works very well for cool stars which have thousands of lines, but hot stars have only a few tens of lines at most, which strongly limits the advantage of multiplexing. Another complication is that this Zeeman technique only measures the projected component of the field in the line of sight, and the true field can be 10 times larger.

In this thesis we describe the best effort ever to search for a magnetic field in one of the brightest O stars, but report only an upper limit. We also describe the discovery of the magnetic field in the B1 III star \(\beta\) Cephei, the hottest star found sofar with a significant field, which clearly demonstrates that massive stars can have a magnetic field (the origin of which, however, remains unclear), and that the techniques are available, but that apparently the sensitivity of the instrumentation is not yet high enough for the application to O stars.

From the above it is clear that the problem of cyclical wind variability can only be solved with a large observational effort. In the last 10 years a number of multisite campaigns has been conducted, organized from Amsterdam (principal investigator H.F. Henrichs), and included space-
and ground-based observatories to study a few carefully selected bright O stars. This thesis concentrates mainly on the analysis of the acquired data on the 4th magnitude O7.5 III star Ξ Persei, which is the most suitable target for this purpose on the Northern Hemisphere. Such campaigns not only involve the logistics of proposal approval and schedule synchronization of sometimes a dozen observatories (all with their own procedures), but also the non-trivial and time-consuming task of reducing the data from equally many different sources. The reduction has to be done in one place to ensure the required homogeneity, which appeared essential for a sound Fourier analysis. Several new techniques have been developed, in particular to extract quantitative information on periodicities with their uncertainties. This analysis will therefore serve as a case study for the further analysis of other O stars, for which data has been obtained.

In this introduction we outline the importance of the study of massive stars, and review the present insights on stellar winds and their variability. We conclude with a summary of this thesis.

1.2. The role of O and B stars

The most massive stars, with spectral type O and B, and masses ranging from 8 to 100 M☉, play an important role in the universe. First of all they are hot and luminous with temperatures between 20,000 and 50,000 K and luminosities ranging from 10⁵ to 10⁶ L☉. With these luminosities they contribute to a major part of a galaxy’s luminosity. The brightest supergiants are visible as individual stars in other galaxies up to the Virgo cluster (at 20 Mpc). These stars have a strong wind, which is driven by the radiation pressure of the star. Mass-loss rates of 10⁻⁶ to 10⁻⁵ M☉yr⁻¹ are typical for O and early B stars. This material is expelled with velocities up to 3000 km s⁻¹, which makes them very important contributors to the energy and momentum budget of the interstellar medium. This mass loss also has a large impact on their evolution. The most luminous stars can lose half of their initial mass during their lifetime. After various stages of evolution the ongoing mass loss finally leads to the exposure of the hot stellar core when the star becomes a Wolf-Rayet star. Ultimately, these stars end their active lives with a violent supernova explosion and enrich the interstellar medium with heavy elements. The shock waves from these explosions trigger new star formation, whereas the remnants, in the form of neutron stars and black holes, are responsible for many energetic phenomena in the universe.

In order to understand the evolution of massive stars knowledge about mass loss and stellar winds is essential. It is therefore important to study in detail the mechanisms at work in radiatively driven stellar winds. Similar physical processes act in other outflows like in quasi-stellar objects (QSOs) and active galactic nuclei (AGNs). Like in many other astrophysical problems, the observed variability provides an important diagnostic for the study of these processes, deriving the relevant physical parameters and to test current theories.

1.3. Radiatively driven stellar winds: a brief history

Massive stars were long suspected to have dense stellar winds driven by the radiation pressure of the star. Milne (1924) predicted the high-speed ejection of atoms from stars and noted the instability of line-driven outflows and their susceptibility to line-shocking (see the review of Lucy 1998). In the C III λ5696 emission line of the O9.5 Ia star α Cam Wilson (1938) detected wings which extend up to 1500 km s⁻¹, and interpreted this as being due to high-speed outflow.

The most important evidence for hot star winds was provided by the discovery of P-Cygni type profiles in the UV resonance lines of C IV, N V and Si IV in the spectra of O and B supergiants obtained with rocket experiments by Morton (1967). These profiles are characterized by a blue-shifted absorption part convolved with an emission profile centered at rest wavelength. The blue-shifted absorption trough is formed by material in front of the star flowing towards the observer, whereas the emission part originates in the outflowing gas all around the star. The most negative velocity at which still absorption is found (called the velocity edge) is identified with the terminal velocity (v∞) of the wind. Since the observed v∞ is several times larger than the escape velocity (vesc), it became immediately clear that these stars are losing mass. Abbott (1982) found that the mass-loss rate M ∝ L^1.6 and that v∞ scales with vesc. This gave strong support to the idea that radiation pressure provides the driving force.

It can easily be shown that electron (Thomson) scattering alone is not sufficient for providing the radiative acceleration. Lucy & Solomon (1970) showed that the line opacity from strong lines formed in the wind makes an important contribution to the driving of stellar winds. Castor, Abbott & Klein (1975, CAK) further developed the radiation-driven wind theory by including the contribution of weak lines. They provided a theoretical basis for the observed relation between M and L and v∞ and vesc. Integration of the equation of motion results in the widely used velocity law for the relation between velocity v and distance r to the center of the star, which takes the form

\[ v(r) = v_\infty \left(1 - \frac{R_*}{r}\right)^{\beta} \]  

(1.1)

According to the CAK-theory \( \beta = 0.5 \). Later improvements (Friend & Abbott 1986, Pauldrach et al. 1986) resulted in \( \beta = 0.8 \), which is in good agreement with observations. Also the predicted mass-loss rates and terminal velocities are in fair agreement with the observations. Furthermore, a comparison between observed and predicted line profiles reveals fundamental parameters such as mass, radius, and luminosity of the star (see Kudritzki & Hummer 1990 and Kudritzki 1998 for reviews). Under certain
**Introduction and summary**

**Fig. 1.1.** Dynamic quotient spectra obtained with the IUE satellite of UV resonance lines of 10 Lac O9 V (N V, $v \sin i=31$ km s$^{-1}$), 19 Cep O9.5 Ib (Si IV, $v \sin i=75$ km s$^{-1}$) and 68 Cyg O7.5 III:n(f) (Si IV, $v \sin i=274$ km s$^{-1}$), showing the difference in progression of the DACs in slow and a fast rotators (from Kaper et al. 1996). The upper panel shows an overplot of the average P Cygni profile (thin line) and the amplitude of the variability. The middle panel contains an overplot of the spectra. The lower panel is a grayscale representation of quotient spectra (obtained after division by a template) with time running upwards, where the same scale was used in the three figures. The rest wavelengths of the doublets are indicated with vertical tickmarks. The horizontal velocity scale is relative to the strongest line of the doublet, corrected for the radial velocity of the star.

conditions these fundamental parameters can also be derived from H$\alpha$ line profiles (Puls et al. 1996, Petrenz & Puls 1996). Since H$\alpha$ can even be observed in O stars in external galaxies they provide an excellent standard candle for distance measurements (Kudritzki 1998).

### 1.4. Wind variability

After the pioneering work by Morton (1967) a systematic study of UV spectra was initiated with the launch of the *Copernicus* satellite. In the ultraviolet P Cygni profiles of many OB stars unexpected features were discovered, the so called high-velocity narrow absorption components (e.g. Underhill 1975, Morton 1976, Snow & Morton 1976, Lamers et al. 1982), of which the nature was totally unclear, although they were obviously formed in the wind. With the launch of the *International Ultraviolet Explorer* (IUE) satellite in 1978, operated jointly by NASA from Goddard Space Flight Center in Greenbelt, and by ESA from Villafrance near Madrid, a new very fruitful era of UV spectroscopy started. It soon appeared that these narrow absorption components could vary in strength and shape on timescales of a few hours and that they occurred in the majority of OB and Be stars (Henrichs 1984). Throughout the operational lifetime of the IUE satellite (1978 – 1996) the study of variability of these UV lines has been one of the focal points and several thousands of UV spectra of OB stars have been recorded. From the obtained time series it became clear that the variability is not chaotic, but occurs in regular patterns (e.g. Pinjia et al. 1987, Henrichs et al. 1988). Broad absorption features appear at low or intermediate ($0.2-0.5 v_{\infty}$) line-of-sight velocity and narrow in width when they approach the terminal velocity. Because of their distinct appearance these features were since then called *discrete absorption components* (DACs).
At first sight, such a strong variability is not expected in radiation-driven winds, in which the dynamics is determined by the \( \sim \) constant luminosity and gravity of the star. However, a line-driven wind is expected to show random fluctuations due to a potent instability: a small-scale increase in the flow speed Doppler-shifts the local line frequency out of the absorption shadow of the underlying material, leading to an increased radiative force which then tends to further increase the flow speed (Lucy & Solomon 1970, Owocci & Rybicki 1984). However, these fluctuations most likely lead to small-scale variability and not to the large-scale regular DAC variability. The black troughs in UV resonance lines and the observed X-ray emission from O stars are, however, well explained by this kind of instability (Lucy 1982, Owocci et al. 1988, Puls et al. 1993).

As first suggested by Prinjia (1988) and Henrichs et al. (1988) and later confirmed by other studies, in particular by the extensive study of the DAC properties in ten different O stars (Kaper et al. 1996, 1999a), it was found that the DAC variability is cyclic and that the recurrence times are proportional to the reciprocal of the projected rotational velocity (vsini). Figure 1.1 shows some examples of developing DACs in progressively faster rotating stars. This strongly suggests that the variability is caused by some structure which corotates with the star. The possibility of the presence of Corotating Interaction Regions (CIRs) in hot-star winds, in analogy to what is observed in the inhomogeneous solar wind was first proposed by Mullan (1984, 1986). In this model the emergent wind flow is perturbed by some inhomogeneity at the surface of the star, which causes an azimuthal asymmetry. Such a structure produces a local change in the wind flow properties. Further-out in the wind, the slow material is caught up by faster wind coming from below and collides, thus forming shock-fronts which corotate with the star. Prinjia & Howarth (1988) suggested CIR-like structures to explain the DAC behavior in the O7.5 III star 68 Cyg. Cranmer & Owocci (1996) performed the first hydrodynamical computations of a CIR model (see Fig. 1.2), and were able to reproduce the behavior of the DACs. They did not specify the nature of the perturbation at the bottom of the flow. One could think of non-radial pulsations (which divide the star in oppositely moving sectors) or magnetic field configurations (with different temperature and brightness).

1.5. Observational strategies

A number of extensive timeseries of UV spectra of OB stars have been obtained, showing the detailed evolution of DACs, following a pattern typical for a given star (e.g. Massa et al. 1995 presenting the IUE MEGA campaign including \( \zeta \) Pup, HD 64760 and HD 56980), Kaper et al. 1996, 1999a), but these observations gave no clue about the origin of the DACs. For this purpose one needs to study the wind structure as a function of distance to the star, which requires elaborate multiwavelength observations. In Fig. 1.3 a schematic map of the line-forming regions in the wind is given. The UV resonance lines (Si IV, N V and C IV) are formed throughout the wind, while e.g. the sub-ordinate lines of N IV and H\( \alpha \) are formed closer to the star. Several other optical lines are formed in the photosphere and in the transition region. The photospheric lines are used to study the stellar surface, e.g. for the presence of non-radial pulsations. Polarization measurements to search for magnetic fields can also be done using optical lines.

The first campaign to study the O7.5 III star \( \xi \) Per, the main target of this thesis, was organized in 1989, which included IUE observations with ground-based support from the Calar Alto Observatory in Spain and Kitt Peak. This yielded the insight that the variability of the wind is already detected near the stellar surface (Henrichs et al. 1994). Furthermore, the pulsational properties of the star could be determined (de Jong 2000b, Chapter 5). The latter was only possible after many modeling efforts to describe and anal-
use the pulsational properties of rotating stars. More co-ordinated optical and UV observations were carried out in 1991 (Kaper et al. 1997), including Hα measurements. This study showed that the wind variability by DACs is also reflected by the variations in the Hα line. In October 1994 an extensive campaign of 10 days of IUE observations and simultaneous multi-site ground-based observations of ξ Per (and a few other O stars) was organized. This star has a DAC period of 2 days. This campaign provided a wealth of information about the wind structure (de Jong 2000a, Chapter 2). All the UV lines and the Hα line show a periodicity of 2.09 d and are strongly correlated. We could only explain these observations by the presence of multiple CIRs and a stellar rotation period of 4.18 d. This is remarkable because the implied radius of the star, given the observed projected rotation velocity, has to be in this case at least 17 R☉, which is rather large for a luminosity class III giant (Kaper et al. 1999a, de Jong et al. 2000a, Chapter 2).

Model atmosphere calculations should show whether this conclusion is justified.

Since the IUE satellite was switched off in 1996 (because of budgetary reasons), only the optical lines (especially Hα) can be observed for further studies of O star winds. In principle the Hubble Space Telescope could be used for this purpose as well, but in practice it is not feasible to obtain continuous observations with the HST for a number of days. In order to get more insight in the optical line profile variability we conducted a new multi-site (MuSiCoS, “MultiSite Continuous Spectroscopy”) campaign on ξ Per in 1996 (Henrichs et al. 1998b, Chapter 3). This time all telescopes were equipped with échelle spectrographs which covered the whole optical range. Unfortunately the IUE satellite was taken out of service just before the start of the campaign (although observing time was allocated). During these campaigns an attempt to measure the magnetic field was made, but only upper limits could be obtained.

1.6. A search for the origin of wind variability

Although the CIR model is now widely accepted as the proper model for the UV wind variability, the origin of the perturbations is still unknown. As described above, non-radial pulsations or magnetic fields are both still valid options, the latter of which still being the strongest candidate.

Several B stars are known to have a magnetic field. The He peculiar B stars are all thought to have magnetically confined winds (e.g. Shore & Brown 1990). Also some β Cep stars, which share their position in the HR diagram overlaps with the He-strong stars were suspected to have a magnetic field because the phenomenology of the stellar wind variations is similar to that of the Bp stars as far as the UV is concerned (Henrichs et al. 1993). We report in Chapter 5 the detection of a magnetic field of 90 G in the slowly rotating star β Cep (B1 IIe), which is modulated with the stellar rotation period of 12 days. This is the hottest upper main-sequence star for which a magnetic field has been dis-covered so far. The field in this star is very weak (the second weakest among all B and A stars with a detected field), but apparently strong enough to control the wind up to 10 stellar radii. For most O stars it is not likely that their winds are magnetically controlled, which also constrains the strength of the field. As an interesting conclusion regarding β Cep we note that this slowly rotating star, which shows intermittent Hα emission similar to the enigmatic Be stars (Mathias et al. 1991, Kaper & Mathias 1995), provides an example in which the Hα emission is not due to rapid rotation such as in other Be stars, but rather due to the presence of a strong enough magnetic field.

It is therefore tantalizing that no magnetic fields have been detected yet in O stars. The Ω 7 V star θ1 Ori C is strongly suspected to have a magnetic field, since the phenomenology of the variations is very similar to that of magnetic rotators of the Ap and Bp group (Stahl 1998). Donati & Wade (1999b) and Mathys (1999) attempted to measure the field of θ1 Ori C, but only found an upper limit of 250 G in the longitudinal component, which means that the polar field strength can be up to 2.6 KG. We made several attempts to measure the field of the Ω 7.5 III star ξ Persei (October 1994 campaign: Henrichs et al. 1998a, de Jong et al. 2000a, Chapter 2; November 1996 MUSICO: Henrichs et al. 1998b, Chapter 3; December 1998 Pic du Midi: de Jong et al. 2000b, Chapter 5), but our best result is an upper limit of 47 G on the disk-averaged longitudinal component of the field. From this we estimate an upper limit of 400-500 G on the polar field strength (de Jong et al. 2000b, Chapter 5). We predict that with at least 4m-class telescopes and very good weather conditions a field detection should be possible with the current instrumentation.

The new generation of polarimetric instruments developed for spectrographs like ESPADONS should make the detection of apparently weak fields possible.

The second candidate for the surface perturbations is the presence of non-radial pulsations (NRP) in which neighboring segments of the star oscillate in different phases. A NRP mode is determined by the parameters ℓ and m. The value of ℓ indicates the total number of node lines on the surface of which m node lines cross the equator. There are three types of NRP: sectoral (ℓ = |m|), zonal (m = 0) and tesseral (ℓ > |m|, m ≠ 0). Examples of these NRP types are shown in Fig. 1.4.

NRP will cause velocity and density perturbations at the base of the wind, but the velocity with which NRP bumps move with respect to the stellar surface depends on the NRP mode and frequency. This velocity has to be added to the rotation velocity in order to derive the timescale it takes for a feature on the surface to cross the visible disk of the star. Until now NRP have been confirmed in six O stars and are suspected in two more (see Henrichs 1999b for an overview). We report in Chapter 4 the discovery of NRPs in the O stars ξ Per (P=3.45 h, ℓ=3) and Λ Cep (P=12.3 h, ℓ=5 and P=6.6 h, ℓ=5). The NRP in ξ Per can most probably not account for the DAC period of 2.09 d, because its pattern.
Fig. 1.4. Examples of non-radial pulsation patterns for selected modes. From the left to the right, the figure shows one sectoral mode ($\ell = |m|$), two tesseral modes ($\ell > |m|, m \neq 0$) and one zonal mode ($m = 0$). For each mode one sees three plots of the radial component of the surface velocity field, is drawn for inclination angles of $i = 90^\circ$, $i = 45^\circ$ and $i = 0^\circ$.

speed and time scale are incompatible. However, other variability in the wind, visible in Hα on time scales of hours, could well be due to NRP (de Jong et al. 2000b, Chapter 5). Beating effects between multiple NRP modes has shown in some Be stars to provide a mechanism to enhance the mass-loss rate at certain pulsation phases (Rivinius et al. 1998). Whether this can give rise to wind variability in O stars is not known. Further studies are obviously needed.

1.7. Towards modeling the wind structure

Extensive time series were also obtained of the Hα line, which maps the inner part of the wind, of most (~40) bright O stars during several years (Kaper et al. 1998). Especially the O supergiants show very complicated line profile variations, which also differ very much from star to star. Main sequence stars do not show such variations in the Hα line, probably because their wind is too weak. This variability is cyclical and is also most likely related to corotating structures. No good model is currently available to explain this behavior in detail. As a first attempt to understand these variations we modeled the Hα variability using a 2D kinematic model in which spiral like structures are accounted for (de Jong et al. 2000c, Chapter 7). We developed a code based on genetic algorithms to search for the best parameters to describe this structure. Due to the limitations of the model only a qualitative agreement could be obtained. This is an essential first step in trying to understand what happens close to the stellar surface in the transition region between the photosphere and the wind.

1.8. Contents of this thesis

In the next three chapters we describe the results of three major campaigns on ξ Persei. Each campaign has its own focus. The first one in October 1994 was the most extensive global campaign, which included UV timeseries of spectra (Chapter 2). We concentrated on obtaining a long period of observations to study the phase relation between the different spectral lines. In the MuSiCoS campaign in November 1996 (Chapter 3) we searched for magnetic fields (simultaneously with line variability) with better instrumentation, but this campaign was hampered by bad weather. The third describes the discovery of non-radial pulsations in the O stars ξ Persei and λ Cephei, based on data obtained in 1989. In Chapter 5 we present the results of the best attempt so far to search for a magnetic field in ξ Per, using the MuSiCoS polarimeter at Pic du Midi in France. Chapter 6 describes
the discovery of the magnetic field in $\beta$ Cephei. In the last chapter we present the results of extensive modeling of the inner part of the wind by means of 2D calculations.

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Chapter 2

A search for the cause of cyclical wind variability in O stars

Simultaneous UV and optical observations including magnetic field measurements of the O7.5III star ξ Persei


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Abstract. We present the results of an extensive observing campaign on the O7.5 III star ξ Persei. This star was monitored continuously in October 1994 during 10 days at ultraviolet and visual wavelengths. The UV observations were obtained with the International Ultraviolet Explorer (IUE). The ground-based optical observations include magnetic field measurements, Hα and He I λ6678 spectra, and was partially covered by photometry and polarimetry. The aim of this campaign was to search for the origin of the cyclical wind variability in this star. We determined a very accurate period of 2.086(2) d in the UV resonance lines of Si IV and N IV and in the Hα line profiles. The epochs of maximum absorption in the UV resonance lines due to discrete absorption components (DACs) coincide in phase with the maxima in blue-shifted Hα absorption. This implies that the periodic variability originates close to the stellar surface. The phase-velocity relation shows a maximum at −1400 km s^{-1}. The general trend of these observations can be well explained by the Corotating Interaction Region (CIR) model. In this model the wind is perturbed by one or more fixed patches on the stellar surface, which are most probably due to small magnetic field structures. Our magnetic field measurements gave, however, only a null-decision with a 1σ errorbar of 70 G in the longitudinal component. Some observations are more difficult to fit into this picture. The 2 day period is not detected in the photospheric/transition region line He I λ6678. The dynamic spectrum of this line shows a pattern indicating the presence of non-radial pulsation (NRP), consistent with the previously reported period of 3.5 h. The edge variability around −2300 km s^{-1} in the saturated wind lines of CIV and NV is nearly identical to the edge variability in the unsaturated Si IV line, supporting the view that this type of variability is also due to the moving DACs. A detailed analysis using Fourier reconstructions reveals that each DAC actually consists of 2 different components: a ‘fast’ and a ‘slow’ one which merge at higher velocities.

2.1. Introduction

All O and many B stars show cyclical variability in their stellar winds. The most prominent features are the Discrete Absorption Components (DACs) observed in UV P Cygni profiles. They accelerate through the blue-shifted absorption trough in a few days until they approach the terminal wind velocity (v∞). The saturated profiles, in which DACs cannot be observed, show regular shifts in their blue edges, up to 10% in velocity. A key observation is that DACs appear periodically (Henrichs et al. 1988, Prinja 1988), and that their recurrence timescale is shorter in stars with a higher (projected) rotational velocity. This strongly suggests that stellar rotation determines the timescale of variability in OB-star winds. For a more extensive introduction and various examples of DAC behavior the reader is referred to Kaper et al. (1996).

The main unsolved issue is where the DACs originate. A comparison of different time series (see Kaper et al. 1999a) of DACs in the Si IV profile shows that for a given star the same recurrence timescale is observed over many years, but detailed changes occur from year to year.
The wind structures can be traced back to very low velocities: basically down to the $v_{\text{sin}i}$ value of the star (see for $\xi$ Per Henrichs et al. 1994). Together with the observed rotational modulation this argues in favor of a model invoking corotating structures which originate at or near the surface of the star. Such models have been proposed by several authors. Underhill & Fahey (1984) investigated the possible effects of magnetic surface structures on O star winds. Mul lan (1984, 1986) proposed that DACs are caused by corotating interaction regions (CIRs) in analogy with the solar wind. On the basis of the CIR model Cranmer & Owocki (1996) performed radiative hydrodynamical calculations to investigate the impact of bright or dark spots on the surface of a rotating star on the dynamical structure of the stellar wind. Such a spot causes a local change in the radial flow properties, for example a higher density and a smaller outflow velocity. Due to the stellar rotation the perturbed stream is curved, so that further-out in the wind, the slow (or fast) stream interacts with the ambient flow resulting in a spiral shaped shock-front that corotates with the star. The physical origin of such a spot is not specified. One could think of the changes in surface temperature and velocity associated with non-radial pulsations (NRP) or a local increase in mass-loss rate due to the presence of a surface magnetic field. Cranmer & Owocki showed that the resulting wind structure gives rise to features resembling DACs in UV resonance lines.

Figure 2.1 symbolically maps out the regions where the most important line regions in O stars are formed. The UV resonance lines are sensitive to (density)$^1$, and are therefore probing the wind over a much larger extent than e.g. H$\alpha$ and the subordinate N\ IV $\lambda 1718$, which fall off as (density)$^2$; the latter, therefore, preferentially sample the innermost part of the wind. The deep photospheric lines are used to study pulsation behavior by means of Doppler imaging techniques. The timescales of pulsation, rotation and wind flow are all in the order of a day, which makes it particularly difficult to disentangle these effects, and a global network of observatories is needed to obtain the required time sampling.

We have conducted several multisite campaigns to simultaneously probe the stellar wind, its base and the underlying photosphere. Here we present the results obtained in the October 1994 campaign on one of the best studied O stars so far: $\xi$ Per O7.5 III(n)((f)). Its bright visual magnitude ($V=4.0$) and the short timescale of the wind variability (Prinja et al., 1987) make this star an excellent target for a detailed study.

2.2. Observations and data reduction

A schematic overview of the campaign is presented in Fig. 2.2. During about 10 days we collected 70 ultraviolet spectra with the International Ultraviolet Explorer (IUE), covering 1150 - 1900Å. Six average magnetic field values, each consisting of 14 measurements, were obtained with the Landstreet UWO Balmer line (H/3) polarimeter attached to the CFHT (see Table 2.2 for the meaning of the acronyms used throughout this paper). High quality H$\alpha$ spectra were collected from various locations around the globe: Canada (DAO, Victoria), USA (Ritter Observatory, Toledo, and BMO, Pennsylvania), Europe (OHP, France and JKT, Canary Islands), and China (BAO), totalling 368 usable spectra covering about 70% per day during the 10-day campaign. The spectra from JKT, OHP and BMO also contain the weak He I $\lambda 6678$ absorption line. Average sampling times were 3 hours for the UV spectra and a half an hour for the optical spectra. Spectropolarimetry was conducted at Pine Bluff Observatory (Wisconsin, USA, observer Bjorkman), but only one spectrum could be obtained. In this section we describe the basic data reduction procedures. Further processing for ensuring a homogeneous dataset is described in Sect. 2.3.1.

2.2.1. Ultraviolet spectroscopy

High-dispersion ultraviolet spectra ($R \approx 10,000$) were obtained with the Short Wavelength Prime (SWP) camera on board the IUE satellite (Nichols, Henrichs, Scheers). The
log of these observations is presented in Table 2.1. For a detailed description of the data reduction we refer to Kaper et al. (1996). We used the Starlink IUEDR software package (Giddings 1981). Interstellar lines were used to align the wavelength calibration. The echelle-ripple correction was performed with the method described by Barker (1984). Reseau marks were removed by linear interpolation. The spectra were mapped on a uniform wavelength grid of 0.1 Å. The signal-to-noise ratio of the spectra is about 30 (see Henrichs et al., 1994).

2.2.2. Magnetic field measurements

Magnetic field measurements were obtained with the University of Western Ontario photo-electric Pockels cell polarimeter attached to the 3.6m Canada-France-Hawaii Telescope at Mauna Kea (observers Bohlender, Hill). With this instrument the Zeeman-splitting due to the longitudinal component of a magnetic field is measured in the Hβ line profile. The red and blue components can be discerned, because they are circularly polarized in opposite directions. The 2-channel photo-electric polarimeter uses two narrow band (5 Å) filters to measure the red and blue wings of the Hβ line, each centered a few Å from the line center at the steepest gradient of the line to ensure the maximum signal (Landstreet 1982 and references therein).

2.2.3. Optical spectra

Long-slit spectra centered on Hα were obtained at DAO, JKT, OHP, TO, and XO whereas échelle spectra were obtained at BMO and RO. The ESO MIDAS package was used for the reduction of BMAO, DAO, OHP and RO data. The spectra from JKT were reduced using NOAO’s IRAF package. A large effort went into the reduction of the data to ensure sufficient homogeneity for an accurate analysis. We describe for each observatory the specific reduction strategy, methods and problems.

(i) Observatoire de Haute Provence

Hα time series were obtained at OHP (observer Kaper) at the coude focus of the 1.52m telescope. We used the Aurélie spectrograph with the 2 × 2048 Thomson CCD detector and grating #7 to obtain high-resolution spectra (R=70,000) in a wavelength region from 6500 to 6620 Å. During all runs, calibration frames were obtained regularly in order to correct for the bias and dark current levels of the CCD. Th-Ar and tungsten flat-field exposures taken at ~2 hour intervals through the night provided wavelength calibration and correction for pixel-to-pixel variations of the detector, respectively. The exposure times were about 10 minutes in which a S/N of 300 was obtained.

(ii) ING La Palma, JKT

Hα and He I λ6678 spectra were obtained using the RBS spectrograph on the 1.0m Jacobus Kapteijn Telescope.
at the Roque de los Muchachos observatory on La Palma (observer Telting). We used the 2400 grating with a resolution of 15,000 covering the wavelength range from 6450 to 6650 Å. Cu-Ne calibration frames were taken after every movement of the telescope, since the spectograph is attached to it. A S/N of 300 was achieved in exposure times of 5 minutes.

(iii) Black Moshannon Observatory

The spectra were obtained with the 1.6m telescope of BMO (State College, Pennsylvania) using a fiber-fed echelle spectograph and a Texas Instruments 432 × 808 CCD with 15μ pixel size (observers Neff and O’Neal). They contain 15 orders covering a total wavelength range from 4780 to 8990 Å, except for a number of gaps between the orders. We first subtracted an average bias from each frame. Secondly, the positions of the orders were located using the Hough method (see Ballester, 1994). The background was fitted by means of a 2D polynomial fit over the inter-order space and subtracted.

The orders were extracted using the optimal extraction method (see Horne, 1986). The orders were flatfield corrected after the extraction procedure. We did not apply flatfielding to the original frames because the orders were too narrow to remove the intrinsic light variations in the original flat fields. The wavelength calibration was done by means of a 2D polynomial fit on the positions of the Th-Ar lines which were constrained by the echelle relation between the absolute order number and the central wavelength. Hα lies close to an order edge which made the normalization of this line rather difficult.

(iv) Dominion Astrophysical Observatory

At DAO (observer Jiang) we obtained long-slit spectra ranging from 6350 to 6850 Å with the spectograph including a Reticon detector attached to the 1.2m McKellar telescope in the couéd focus. The dispersion was 10 Å mm⁻¹. The spectra contain 20 rows of starlight and no background. We had to correct the flat fields for non-uniform illumination by smoothing them row by row over 250 pixels in the direction of the dispersion. Secondly, we divided the flat fields by the smoothed versions, which resulted in frames containing only pixel to pixel variation. We averaged the bias frames per night and performed the standard bias and flat-field calibration. The final spectra were obtained with the optimal extraction method and with a wavelength calibration using Th-Ar spectra.

(v) Xinglong Observatory

At XO (observer Cao) we obtained low dispersion (R=2500) long-slit spectra ranging from 5691 to 6928 Å with the Cassegrain spectograph attached to the 2.16m telescope, using the maximum available dispersion of 50 Å mm⁻¹. The spectra consist of 70 rows of starlight and sky background. The rest of the CCD below and above the stellar slit was used for the Ne lamp for wavelength calibration. Per night all bias frames and flat-field frames were averaged. A bias frame was subtracted from the raw science frames and flatfields. The science frames were afterwards divided by appropriate flatfields.

The Ne-lamp spectra were only taken at beginning and the end of the night, so we could not use them to correct for the wavelength changes during the night. Furthermore, they were taken through a different slit which may have caused additional errors in the line positions. The spectra did indeed shift during the nights. We used a Diffuse Interstellar Band (DIB) at 6612 Å to determine and correct for the shift of each spectrum.

(vi) Ritter Observatory, Toledo USA

The 1m Ritter telescope (observers Gordon and Morrison) has a Cassegrain-mounted echelle spectograph with a maximum resolution of 60,000. The detector was an EEV CCD chip with 1200 × 800 pixels of 22.5μ size.

The reduction of the data was done in the Interactive Data Language (IDL) with a specialized program written for Ritter Observatory data (Gordon 1995) based on methods detailed in Hall et al. (1994). The average bias and flat field were constructed on a pixel-by-pixel basis allowing the removal of cosmic-ray hits. The average bias was subtracted from the average flat field, objects, and wavelength calibration frames. The flat field was used to determine the order and background templates. The background template was used to remove the scattered
light from the flat field, objects, and wavelength calibration frames after fitting a polynomial to the interorder background on a column by column basis. Cosmic ray hits in the objects and wavelength calibration frames were removed by comparing them with the average flat field. The orders were extracted by using a profile-weighted extraction method. The wavelength calibration was accomplished by scaling all the comparison lines to one order and fitting a polynomial to the result iteratively removing points until a preset standard deviation was achieved.

(vii) Tartu Observatory

At TO in Estonia (observer Kolka) we obtained 54 intermediate resolution (R=16,000) Hα spectra using a long-slit Cassegrain spectrograph mounted on the 1.5m telescope. A ST-6 CCD camera was used and Ne arc spectra were taken for the wavelength calibration. In all cases the exposure time was 10 minutes, yielding a typical S/N of 60–80. In order to increase the S/N to match the other observatories we averaged blocks of spectra within time intervals of one hour (24 averaged spectra). The wavelength calibration was very problematic, because only two Ne lines fell within the short wavelength range of the detector. An attempt to find the non-linear dispersion coefficients was made by comparing the ξ Per spectra with simultaneously taken spectra from other observatories, but no reliable match could be achieved, and consequently we could not use these spectra for the time series analysis.

2.2.4. Removal of telluric lines

Many of the Hα line profiles are heavily contaminated with about 20 distinct telluric lines. This contamination varies from night to night, mainly as a function of the amount of water vapor in the lower atmosphere. To remove these effects we first created a list of the telluric lines from a spectrum of a star with very broad lines (co-observed with the target star) and constructed a template telluric spectrum, which can be scaled as a function of average line strength and line width, and shifted in wavelength. In an automatic fitting procedure the contaminated source spectrum is divided by such a template and the three parameters are adjusted until the telluric lines are removed. Here follows a description of the method.

1. The telluric lines were identified by eye in the spectral region around the Hα profile of ζ Aql, A0V, which has strongly Stark-broadened Balmer lines. A spline curve was fitted through selected uncontaminated points and the stellar spectrum was divided out, which leaves a spectrum S(λ) of only telluric lines. This spectrum was inverted (S′(λ) = 1 − S(λ)) and modeled using a multiple Gaussian function:

\[ S′(λ) = \sum_{i=1}^{N} A_i \exp \left[-\left(\frac{λ - λ_{0i}}{w_i}\right)^2\right] \]  \hspace{1cm} (2.1)

\[ T(λ) = \sum_{i=1}^{N} (1 - e^{-τ_i Δτ}) g(λ_i), \]  \hspace{1cm} (2.2)

\[ g(λ) = \exp \left[-\left(\frac{λ - λ_{0i} - Δλ}{w_i Δλ}\right)^2\right] \]

\[ τ_i = -\ln(1 - A_i) \]

3. A region containing strong telluric lines is chosen in the source spectrum. In this region a polynomial P(λ) is fitted to the continuum, which is compared to a residual spectrum R(λ) = S(λ)/ (1 − T(λ)). When the residual contains the least amount of telluric lines the following χ² should be minimal:

\[ χ² = (R(λ) - P(λ))² \]  \hspace{1cm} (2.3)

4. We adjusted Δτ, Δw and Δλ and iterated from step 2 until a minimal χ² was achieved using the Fletcher-Reeves-Polak-Ribiere minimization algorithm (see Press et al. 1992).

It was sufficient to make a line list from only a single ζ Aql spectrum of OHP (highest spectral resolution) in order to remove the telluric lines completely from all spectra regardless of observatory or time (see Fig. 2.3 for an example).

2.3. Combination of the datasets

2.3.1. Creation of a homogeneous dataset

For each observatory the spectra were normalized, transposed to the stellar rest frame, converted to velocity and
Table 2.3. List of optical and UV lines studied in this paper. The third and fourth columns list the rest wavelengths used to convert to radial velocity units. The last column indicates the order of the polynomial used for the normalization of the continuum.

<table>
<thead>
<tr>
<th>Line</th>
<th>Type of line</th>
<th>( \lambda_0(1) (\text{Å}) )</th>
<th>( \lambda_0(2) (\text{Å}) )</th>
<th>region excluded in normalization (km s(^{-1}))</th>
<th>polynomial order</th>
<th>Formation region</th>
</tr>
</thead>
<tbody>
<tr>
<td>N\textsc{iv}</td>
<td>resonance</td>
<td>1238.821</td>
<td>1242.804</td>
<td>&lt;3046</td>
<td>0*</td>
<td>wind</td>
</tr>
<tr>
<td>Si\textsc{iv}</td>
<td>resonance</td>
<td>1393.755</td>
<td>1402.770</td>
<td>-3000 to 3000</td>
<td>1</td>
<td>wind</td>
</tr>
<tr>
<td>N\textsc{iv}</td>
<td>sub-ordinate</td>
<td>1718.550</td>
<td>1550.772</td>
<td>-3000 to 3000</td>
<td>1</td>
<td>inner wind</td>
</tr>
<tr>
<td>C\textsc{iv}</td>
<td>resonance</td>
<td>1548.187</td>
<td>6562.817</td>
<td>-3550 to 2360</td>
<td>1</td>
<td>wind</td>
</tr>
<tr>
<td>H\alpha</td>
<td>recombination</td>
<td>6678.154</td>
<td>6683.209 (He\textsc{ii})</td>
<td>-1000 to 1000</td>
<td>3</td>
<td>photosphere + inner wind</td>
</tr>
<tr>
<td>He\textsc{i}</td>
<td>**</td>
<td>6678.154</td>
<td>6683.209 (He\textsc{ii})</td>
<td>-500 to 750</td>
<td>3</td>
<td>photosphere</td>
</tr>
</tbody>
</table>

* Only the red side of this line showed a proper continuum.

** Only present in JKT, BMO, DAO and XO spectra.

![Graphs](image_url)

Fig. 2.4. Timeseries of the Si\textsc{iv} resonance line and the subordinate N\textsc{iv} and H\alpha lines of the O7.5 III star \( \xi \) Per in October 1994. Time is running upwards. Intensity is represented in levels of grey. To enhance the contrast the shown spectra are divided by a template. The Si\textsc{iv} line shows a regular pattern of recurring DACs (period about 2 d) with detailed changes from event to event. A clear correlation can be seen between these DAC events and the absorption in N\textsc{iv} and H\alpha. It is also clear that the H\alpha variations are much more complicated.
rebinned to the same grid. The normalization of the UV lines was done with a straight-line fit through the continuum around each studied line individually. The continua of the optical spectra were much more curved and often contained spurious features. Therefore we normalized them by fitting 3rd order polynomials to a sufficiently large range of carefully selected continuum points. In the case of Hα we removed the telluric lines before normalization, because their presence influences the determination of the continuum level. The used continuum regions are given in Table 2.3. The normalization of the BMO Hα spectra was quite problematic, because this line lies close to the edge of an order, leaving little continuum at one side. Also the XO spectra gave problems: because of the low resolution and a nearby diffuse interstellar band (DIB) there were very few continuum points close to the blue side of Hα. In both cases this degraded the accuracy of the polynomial fits. The continuum around the He I λ6678 line was so poorly defined that we divided the spectra by their time average before normalization, in this way many small-scale spectral features were removed, thereby making the normalization more accurate.

The selected lines were converted to velocity scale using the laboratory wavelengths in Table 2.3. In the case of doublets we used the shorter wavelength of the two lines.

All optical spectra were rebinned to a resolution of 15,000 (20 km s$^{-1}$) which is somewhat higher than the resolution of JKT spectra (10,000) but lower than the resolution of most spectra. We deconvolved the XO spectra by a Gaussian profile with a FWHM of $\sqrt{100^2 - 20^2}$ to increase their resolution artificially.

Finally, we divided the individual spectra by their average (optical lines) or a minimum absorption template (UV lines). This UV template is identical to the one used and described in Kaper et al. 1999a. The resulting dynamic quotient spectra of Si IV, N IV and Hα are shown in Fig. 2.4.

**2.3.2. Comparison between different observatories**

We checked for possible systematic differences between spectra obtained at the different observatories by comparing Hα spectra taken at approximately the same time. We could only compare OHP with JKT, BMO with RO and BMO with DAO. In case of TO this comparison could not be applied, since the TO spectra were matched to simultaneously taken data (see Sect. 2.2.3).

From this comparison we found, unexpectedly, that the RO spectra deviated up to 3% from BMO spectra in the velocity range from 0–300 km s$^{-1}$ whereas the BMO and DAO spectra agreed within the noise. This deviation of the RO spectra did practically not affect the period analysis (see Sect. 2.5.5), probably because these spectra were all taken within a single small time interval.

The XO spectra also seem to differ systematically, although we could only check this by comparing their average with the average JKT spectrum. For XO the average spectrum is about 2.5% less deep in the wings. We were not able to find an explanation for this effect, but we note that the deviations were less severe in spectra of another star (19 Cep), which do not show strong variations on a time scale of one day. However, the period analysis showed that including the XO data causes a strong 1-day signal which was not seen in the analysis of any other combination of observatories. We therefore decided to correct for the deviation in the XO data using a spline fit through the quotient of the average JKT and average XO spectra.

We could not include the TO spectra in our further analysis. Even though the wavelength calibrations were corrected using simultaneously taken spectra from other observatories, equivalent-width comparisons showed unexplained differences, which strongly affected the results of the period analysis.

**2.4. Overview of the observed line profile variations**

Figure 2.4 shows the dynamic quotient spectra including the Si IV, N IV and Hα lines transposed to the stellar rest frame. For construction of the quotient Hα spectra we used a plain average profile. In this figure, the Hα spectra were smoothed by taking a running mean with a time sampling of about 2 hours, where the contributing spectra were weighted according to their epoch within the sampling width. The resulting quotient spectra have a S/N of better than 200. The development of five strong Si IV DAC events can be seen; in most cases the events consist of different components with different kinematic properties. The five strongest DAC events also clearly show up in the low-velocity region of the N IV lines, down to $-200$ km s$^{-1}$, i.e. the $v_{\text{sin}i}$ value of the star. They recur every 2 days.

The same holds for the Hα region: the strongest absorption around $-250$ km s$^{-1}$ is in concert with the epoch of enhanced absorption in N IV at intermediate velocities, and the appearance of a DAC in Si IV around $-1500$ km s$^{-1}$. This unambiguously shows that the wind structures responsible for DACs can be traced down to the stellar surface.

In Fig. 2.7 we show the normalized Hα line profiles in order to illustrate how these profiles change when the enhanced blue shifted absorption occurs. The maxima in equivalent width (EW) measured between $[-500, -200]$ km s$^{-1}$ (see Sect. 2.5.5.1) are indicated by elongated solid tickmarks (the dotted tickmarks in between indicate the EW minima). It is clear that most variations occur in the blue wing where the profile is even visible up to $-600$ km s$^{-1}$ during the EW maximum around MBJD=8. The central line depth and red wing remain relatively constant, with exceptions around MBJD=2.6 and MBJD=6.8 where some emission appears in the red wing.

There are also variations of a different type. In Fig. 2.5 the edge variability near $-2600$ km s$^{-1}$ of the saturated P Cygni profiles of of C IV and N V are shown, compared with the Si IV edge. This variability is not periodic, but it is clear that the three edges move in concert. A relation with
the DAC events is suggested when comparing to Fig. 2.9. Some DAC events reach a higher terminal velocity than others. An explanation might be that they influence the edge differently.

The photospheric He I \( \lambda 6678 \) line shows variability on a short timescale. In Fig. 2.6 narrow absorption features are visible (notably around 0.8, 4.6 and 8.6 MJD) which move too fast from blue to red to be associated with rotation. Our sampling rate is too low to determine a period for these features, but a separate investigation of a dataset with much higher time resolution revealed that they correspond to non-radial pulsations in mode \( \ell =3 \), with a frequency of 6.95 c/d (de Jong et al. 1999a, Chapter 4); see also Fig. 2.19.

### 2.5. Quantitative analysis

In the previous section we showed the characteristics of the cyclical wind variability. In the following this variability will be quantified by means of equivalent width measurements, temporal variability spectra (TVS), DAC modeling and timeseries analysis.

#### 2.5.1. Equivalent Width measurements

We measured the equivalent widths (EW) in H\( \alpha \) and the UV resonance lines in velocity intervals where the appearance of DAC and edge variability is most prominent. For the error computation we followed Chalabaev & Maillard (1983)

\[
\frac{F_\lambda}{\sigma_{\text{exp}}} = 29.8 \tanh \left( \frac{F_\lambda}{321} \right),
\]

where \( F_\lambda \) is the flux (in units of flux density numbers), and \( \sigma_{\text{exp}} \) is the expected noise. The results are shown in Fig. 2.8.

The shape of the EW curve is very similar for both the edge and the blue-shifted absorption part of the line profiles of Si IV, N IV and C IV. The pattern of slow increase and sudden decrease in absorption repeats itself about every 2 days. This behavior is also found in the H\( \alpha \) line in the absorption region between \(-500 \) to \(-200 \) km s\(^{-1}\), although
Simultaneous UV and optical observations and magnetic field measurements of the O7.5III star ξ Persei

the absorption maxima lag about 0.2 days behind as compared to Si IV. However, the EW of the Hα and He I lines between -200 and 200 km s^{-1} show clearly a different pattern with a strong decrease in EW around MBJD=7 just half a cycle before the strongest absorption at MBJD=8.

The EWs of the edge of Si IV, C IV and N V behave similarly, although there is no clear variation with a 2-day period. The EW drops quite suddenly after MBJD=6.5. Note also that the lowest EW in all edges coincides with the highest EW in Hα at MBJD=8. If they are physically related this would be remarkable considering the entirely different wind regions which these lines should probe.

2.5.2. Magnetic field analysis

The longitudinal magnetic field strength is proportional to the differential polarization measured between the two wings of the line, which is in our case Hβ (Mathys 1989). The derived values with their 1-σ error bars are listed in Table 2.4 and plotted in Fig. 2.8. Each quoted value is an average of 14 successive measurements. The average value of all measurements is 27 ± 70 G, i.e. consistent with a non-detection. No periodicity or correlation is apparent.

Although this is the most accurate search for a magnetic field in an O star to date, it is clear that the error bars on the magnetic field measurements are still too large to claim a significant field detection. A possible coincidence is that the lowest measured value of $B_\gamma=-79\pm83$ G at MBJD=5.875 occurs near a maximum EW in Si IV (see Fig. 2.8 and Table 2.4), but no specific meaning can be assigned to this without further simultaneous measurements.

2.5.3. DAC modeling

We used a least-absorption template in order to isolate the migrating DACs from the underlying P Cygni profile. This template was derived using a rigorous statistical method
Fig. 2.9. DAC model fit parameters compared to the variability in N IV and Hα. In panel (d) the DAC fits are overplotted on the dynamic spectrum of Si IV. The contour plot in the same panel shows the absorption features in N IV (same IDFT as in Fig. 2.16). The EW of Hα between [−500,−200] km s$^{-1}$ is shown in panel (e). The arrows in panel (d) indicate the times of maximum EW in Hα. In panels (a), (b) and (c) are shown the fit parameters $v_t$ and $\tau_c$ and the column density $N$, respectively. The grey-scale conversion is the same as in Fig. 2.4. The overall hollow curvature in panel (d) is called phase bowing (see Sect. 2.6).

which is described in detail in Kaper et al. (1999a). The N IV and Si IV quotient spectra shown in Fig. 2.4 were obtained with this template.

The DACs in the quotient spectra of Si IV are modeled using the method described by Henrichs et al. (1983, see also Telting & Kaper 1994). The DACs are assumed to be plane-parallel slabs of material in the line of sight. They give rise to a Gaussian shaped absorption component at a velocity $v_c$ with a broadening parameter $v_t$ and central optical depth $\tau_c$. The intensity of the component is described by:

$$I(v) = \exp\left(-\tau_c \exp\left[-\left(\frac{v - v_c}{v_t}\right)^2\right]\right)$$  

(2.5)

We assume that the DACs can be modeled simultaneously by using the doublet separation and the ratio of oscillator strengths of the doublet members (Table 2.3). The 3 free parameters are fitted for each pair of DACs by means of a $\chi^2$-method. In most cases a simultaneous fit with 3 or...
4 DACs is needed. The column density $N_{\text{col}}$ is computed according to Henrichs et al. (1983):

$$N_{\text{col}} = \frac{m_{e}c \sqrt{\pi}}{\pi e^{2}} \frac{\tau_{e}v_{l}}{f \lambda_{0}} \left(1 + v_{e}/c\right)$$

(2.6)

where $\pi e^{2}/m_{e}c$ has its usual meaning, and $f$ is the oscillator strength of the transition at wavelength $\lambda_{0}$. The results of the model fits are shown in Fig. 2.9. The evolution of the column density shows that a strong DAC appears every 2 days. Such DACs are always preceded by a weaker DAC, which peaks in $N_{\text{col}}$ about one day earlier. All DACs become detectable between $-1000$ and $-1500$ km s$^{-1}$ with a $v_{l} \approx 700$ km s$^{-1}$ and subsequently become narrower as they move towards more negative velocities. The strong events coincide with the absorption features in N IV and the maximum EW in the blue wing of Hα. Their column densities alternate between about $4 \times 10^{12}$ cm$^{-2}$ and $6 \times 10^{12}$ cm$^{-2}$. The weak events peak about halfway the strong events. There is no analogous feature in the N IV and Hα lines. We also see a persistent component which remains near the edge velocity between MBJD=4.4 and MBJD=6.0, which seems to be unrelated to any other event. An interesting event is the apparent crossing of two DACs around MBJD=6.2 (this also occurred in other years, see Kaper et al. 1999a).

2.5.4. Temporal Variance Spectra

The amount of variability in the line profiles can be quantified by computing the temporal variance spectra (TVS, see Fullerton et al., 1996). The TVS at velocity bin $j$ is defined as follows:

$$TVS_{j} = \frac{1}{N - 1} \sum_{i=1}^{N} \left(\frac{F_{ij} - \bar{F}_{j}}{\sigma_{ij}}\right)^{2}$$

(2.7)

Here $F_{ij}$ and $\sigma_{ij}$ are the flux and the corresponding errors for wavelength bin $j$ in spectrum $i$. $\bar{F}_{j}$ is the average flux in bin $j$. The $\sigma_{ij}$ values of the UV lines were computed using Eq. 2.4.

For the optical lines we treated each observatory differently, because the dependence of S/N on the flux is closely related to the type of spectograph, quality of the calibration frames and the reduction method. In case of Hα the variability is quite large compared to the noise. We therefore only simply measured the noise in the continuum next to the profile assuming that it is proportional to $\sqrt{F}$. In contrast, the amplitude of the variability in the He I $\lambda 6678$ line is only slightly larger than the noise. For the noise measurements in this line we took into account the flux level of the unnormalized and in case of échelle spectra also unflattened spectra. We measured the S/N in many continuum points and fitted a Poisson function through these points ($F/\sigma_{F} = \sqrt{gF}$). Using these fits we determined the noise in each velocity bin of the normalized spectra.
The resulting TVS computations are shown in Fig. 2.10. This figure shows clearly that the different lines probe the variability in different regions. He I only shows variability within $\pm v\sin i$ whereas the variations are visible at increasing velocity going from Hα, N IV, Si IV, N V to C IV, which shows variability only near $v_\infty$ (this line is saturated at lower velocities).

2.5.5. Period search

We performed a standard CLEAN analysis (Roberts et al. 1987) on all UV and optical lines. In this method the discrete Fourier transformation (DFT) is deconvolved with the DFT of the window function, thereby removing the power which originates in the data sampling. This method gave the least amount of spurious peaks when applied to synthetic data with the same properties as our optical and UV data. Moreover, a comparison between different time series analysis methods by Carbonell et al. (1992) shows that the CLEAN method is most applicable to our randomly gapped and relatively short time series.

2.5.5.1. Periods in EW measurements

To enable a comparison with earlier measurements, our first period search was done on the EW measurements (see Sect.
2.5.1 and Fig. 2.8). The resulting powerdiagrams (see Fig. 2.11) clearly show the same periodicity with a frequency of 0.48 c/d in all UV lines and Hα.

We determined the frequencies of maximum power \(f_0\) by means of Gaussian fits. The errors were derived from the width and height of the peak and the noise level in the powerdiagram. The procedure is described in detail in de Jong et al. (1999a, Chapter 4), whereas the error calculation is based on Schwarzenberg-Czerny (1996).

The resulting values of the period \(P=f_0^{-1}\) and the corresponding 1-\(\sigma\) errors \(\epsilon_P = \epsilon_f/f^2\) are given in Table 2.5. Except for the He i line and the edges of the Si iv, N v and C iv, all periods are the same within a 2-\(\sigma\) confidence limit. The weighted average period of the five similar lines is \(P_1 = 2.087(5)\) d.

Some other periods are also present. The Si iv analysis shows a peak at \(P_2 = 1.021(3)\) d and N iv has within the errors a similar period. In addition, \(P_1\) and \(P_2\) are not present in the EW of Hα and He i between \(-200\) and \(200\) km s\(^{-1}\), and also not in the edges of the UV lines. From the Fourier analysis follows a period of 6.4 days in the edge variability (see Table 2.5). This period probably reflects the gradual change in the edge, since the total length of the campaign was only 10 days.

Sinusoids with the 2.09 d period were fitted through the EWs of Hα, N iv, Si iv, C iv and N v. The times of the maxima are given in Table 2.5. The EW maxima of Hα lag 0.26(5) d behind the maxima of Si iv, but coincide with the EW maxima of C iv and N v. We will later discuss these important phase differences in view of the CIR model.

2.5.5.2. Periods in line profiles

In addition to the period analysis of the integrated line profiles we also carried out such an analysis with the CLEAN method for all the studied lines in each velocity bin. The resulting power is plotted as function of velocity and frequency in the contour plots in Figs. 2.12 and 2.13. These figures show that the 2.09 d period is present in all UV lines and that the distribution of the power as a function of velocity is similar to the signal distribution in the TVS (see Fig. 2.10).

The periods and the corresponding uncertainties were calculated with the same method as in Sect. 2.5.5.1. Data points with a power less than 10% of the maximum power and/or with deviation too large compared to their errors were discarded (Chauvenet criterium, see de Jong et al. 1999a, Chapter 4). The remaining datapoints are plotted in Fig. 2.14. The resulting weighted average frequencies are given in the second part of Table 2.5. The best determined period is the 2.086(2) d period in the Si iv line, which we adopt as the wind period. All other periods in the EW and lvp's agree with this value within 1\(\sigma\) (except for the lvp period in Hα). The difference with the weighted average of the EW period is less than 1\(\sigma\).

As can be seen in Fig. 2.14, the period varies a little as a function of velocity, in particular in Si iv and N iv, where a similar trend can be observed, but in both cases the uncertainties are too large to make it significant. We note in this respect that Fullerton et al. (1997) found a similar trend in the star HD 64760, but could not make it significant either.

2.5.6. Comparison of cyclical behaviour in different lines

In order to search for a connection between the variations in the UV and optical spectra we compared the phases of the periodic modulations. The phase of the signal can be extracted from the CLEANed Discrete Fourier Transform (CDFT) using the following representation:

\[
F(t) = \Delta f \sum^{N}_{i=1} \sqrt{2} P_i \cos(2\pi \left[ f_i(t - \bar{t}) + \phi_i \right])
\]  

(2.8)

Here \(F(t)\) is the flux as a function of time \(t\), \(f_i\) is the frequency, \(\Delta f\) is the frequency interval \(f_{i+1} - f_i\), \(P_i\) is the power, \(\phi_i\) is the phase of component \(i\) and \(\bar{t}\) is the average time of the sample. Hereafter we consider only the phase of the component which has the maximum power (at the periods given in Table 2.5). By convention \(\phi\) is defined between 0 and 1, with \(\phi = 0\) at a maximum in \(F(t)\). Note that features with a larger phase arrive earlier in time. We derived the phases from a multiple least-\(\chi^2\) cosine fit in which the errors were computed using a Monte-Carlo approach (see de Jong et al. 1999a, Chapter 4). The time corresponding to the maximum flux in a component, listed in Table 2.5, is \(t_{\text{max}} = (N - \phi')/f\), with \(N\) an integer value between \(-\infty\) and \(+\infty\) and \(\phi' = \phi - f \bar{t}\) (from Eq. 2.8), which is needed to correct for the different epochs in the datasets.

The phases as a function of velocity derived from the cosine fits are shown in Fig. 2.15. In case of Hα, C iv and N v only the 2.09 d period was fitted. The N iv and Si iv also contain a 1.02 d period which was fitted simultaneously. We consider the phases from the cosine fits more reliable than those derived from the CLEAN algorithm, since the latter values have no errors and the effects of removing the window function are not clear.

In all UV lines the phase is not constant throughout the spectral line, but peaks at a certain velocity: at about \(-1400\) km s\(^{-1}\) for Si iv and N iv. This so-called phase bowing, which is best seen in the Si iv line, is characteristic for spiral-shaped structures in the wind (see e.g. Owocci 1995). We will discuss this further in Sect. 2.6.

We also computed the difference in phase between Si iv and N iv, N iv and Hα, and C iv and Si iv. The average phase difference between Ho and the N iv line in the interval \([-460, -200]\) km s\(^{-1}\) is only 0.005(32) cycles, from which we savely can conclude that absorption events occur simultaneously in these lines. However, the absorption events in N iv lead the events in Si iv by about 0.1 cycle.
Fig. 2.12. Powerdiagrams of Si IV, N IV and Hα. The upper panels show the normalized average spectrum. The contours indicate the power relative to the highest peak in the diagram, which value is given below each panel. The dashed contours show respectively the 1%, 2% and 5% levels, while the continuous lines show the levels from 10% to 100% (9 levels).

Table 2.5. The main periods found in the periodograms. The columns at the left contain the results of a period analysis based on EW measurements over the intervals indicated. Column 5 gives the epochs when the sinusoid has a maximum. The columns at the right give the results based on a separate period analysis in each velocity bin in the listed interval. The last column gives the weighted average of the calculated periods over the interval. The underlined value is our adopted best-determined period.
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in the region between −500 and −1000 km s\(^{-1}\) (see Fig. 2.15).

To investigate the origin of these phase lags we reconstructed the dynamic spectra by means of an inverse DFT (IDFT), based on selected frequencies only, which contain the 2.09 and 1.02 d modulation but no noise frequencies. For Si IV and N IV only the fitted cosines of these two periods were used to reconstruct the dynamic spectra. For Hα the CDFT was inverted using all frequencies from 0.1 to 5 c/d and above the noise threshold. Certainly above −200 km s\(^{-1}\) this line is too complex to describe with a single frequency.

Fig. 2.16 shows the resulting dynamic spectrum. In the figure we overplotted symbols indicating the maximum of the cosine fits, \(t_{\text{max}}\). An interesting phenomenon is visible in this figure. In a close inspection of the Si IV panel one can see that the DACs consist at least of two components at low velocity, in correspondence with the DAC fits (see Sect. 2.5.3 and Fig. 2.9). These components could be interpreted as due to separate DACs with different kinematic properties, which merge at higher velocities, yielding a stronger single DAC recurring each 2.09 d. The 1.02 d period is needed to account for the weak DACs in between. In terms of the CIR model (see below) the different kinematics could correspond to different curvature angles of spiral-shaped structures in the wind.

2.6. Discussion of the cyclical variability and phase-bowing

The 2.09 d period appears in the Hα, N IV, Si IV, N V and C IV lines in the velocity range from −200 to −2420 km s\(^{-1}\), which covers the whole wind from the rotational velocity of the star up to the terminal velocity. The relation between the absorption features responsible for this period is especially well shown in Fig. 2.9 were we superimposed the maximum of the Hα EW between [−200, −500] km s\(^{-1}\), the N IV absorption features and the DACs of Si IV. All these features closely coincide in time, implying that the prominent DACs in Si IV can be traced down to the surface of the star. Fig. 2.9 also shows that the absorption events are bow-shaped. In Si IV the DACs move blue-ward, whereas the absorption components in N IV move red-ward. We have
also seen that the phase of the 2.09 d period is bowed and has a broad peak around $-1400 \text{ km s}^{-1}$. Thus, according to the Fourier analysis, the Si IV absorption is first detected at about $-1400 \text{ km s}^{-1}$ and then splits into a strong blue component (i.e. the accelerating DAC) and a weak red component which moves further to the red. The red component is too faint to be visible in the original spectra, but the effect is clearly seen in the reconstruction shown in Fig. 2.16.

We have to keep in mind that the DAC modeling with Gaussian fits as described in Sect. 2.5.3 is a phenomenological description which does not specify the actual place and shape of the structures in the wind. For this we have to concentrate on the periodic behavior. One of the most important characteristics is the phase bowing, mentioned above, which can be interpreted as a splitting of the periodic absorption features into two components at different velocities. Kaper et al. (1999a) also found that the phase in the Si IV line of ζ Per peaks around $-1400 \text{ km s}^{-1}$. There is no phase-bowing observed in the other stars discussed in this paper, probably because the DACs are too weak at lower velocities to be observed with the IUE satellite. Another star which shows very prominent periodic variations is the B supergiant HD 64760, which was observed during the IUE MEGA campaign (see Massa et al. 1995). In this star the main period is 1.2 d (Prinja et al. 1995). A second period appears at 2.4 d and the phases as a function of velocity peak at $-710 \text{ km s}^{-1}$ for both periods (Fullerton et al. 1997). They also found that a given phase of the modulations occurs progressively later for lines that diagnose higher energy processes, which is in agreement with the phase lag of 0.1 d which we found between N IV and Si IV (see Sect. 2.5.6).

Owocki et al. (1995) ascribed this variability to banana shaped modulations in the spectrum rather than to migrating DACs, because all the velocity bins seem to vary sinusoidally with time. However, if we look more closely to the time variations in each velocity bin of ζ Per we can see that they also vary sinusoidally in all lines (see e.g. the EW measurements in Fig. 2.8). All the variations can be described with just two periods of 2.09 d and 1.02 d. The latter period is needed to accommodate for the weaker DACs in between the strong ones. The phase of the 1.02 d period behaves differently compared to the 2.09 d period, possibly related to a contamination due to the overlap between the two doublet components in Si IV. HD 64760 has a lower $v_{\infty}$=1500 km s$^{-1}$ and therefore less blending in the Si IV line. Taking all the similarities together there is sufficient ground to conclude that the modulations in HD 64760 and the DACs in ζ Per are similar phenomena. There were also DAC events with a period of 9.8 d observed in HD 64760 which show the classical pattern of acceleration (Fullerton et al. 1997). There seems to be no analogy to these
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Fig. 2.14. Measured period per wavelength bin obtained in the Fourier spectra by means of gaussian fits centered at 2.09 d. The dashed lines show the average of all points. Only the fits with a power of at least 5% of the maximum power are shown.

Fig. 2.15. Phases for the Hα, N IV and Si IV lines as defined in Eq. 2.8 and corrected for the different epochs (see Sect. 2.5.6). In the interval [-460, -200] km s⁻¹ the average phase difference between Hα and N IV is only 0.005(32). Note also the phase bowing in the Si IV lines are similar as well. Kaper et al. derived a period of 2.0±0.2 d which is in good agreement with our observations. From this we conclude that whatever causes the DAC events is stable for at least 7 years. The detailed differences between different years are not larger than the differences between subsequent DAC events.

2.6.1. Long-term behavior

Considering the unknown origin of the variability, it is important to investigate the long-term behavior of DACs in ξ Per. The DAC modeling described in Sect. 2.5.3 shows only the blue-ward moving components and gives results similar to the UV observations from 1987, 1988 and 1991 (Kaper et al. 1999a). These observations also show that strong DACs appear about every 2 days with a weak DAC in between. The DAC parameters at these epochs are comparable: $N_{\text{col}}$ varies between 3.6 and $5.5 \times 10^{14}$ cm⁻², which is about the same in our observations. The $v_t$ and $\tau_c$ values

events in ξ Per. In our case only the edge variability of the UV lines could give an indication of a timescale which is longer than our observation run.

Currently, the cyclical wind variations which we described above are best explained by the Corotating Interaction Regions (CIR) model in which the emergent wind flow is perturbed by some structure at the surface of the star. Such a structure causes a local change in the radial flow properties. Due to the stellar rotation such a stream is curved. Further-out in the wind, slow and fast moving streams collide and form a shock-front that corotates with the star. This model was proposed by Mullan (1984) in analogy with the CIRs in the solar wind. Owocki et al. (1995) first showed that spiral shaped corotating structures give a good explanation for the observed bow-shaped isoflux contours in HD 64760. They used a kinematic model in which spiral “streak” lines are rotated through the line of sight. They showed that the part from the spiral at intermediate velocities enters the absorption column first, followed simultaneously by the low- and high-velocity parts. The low-velocity part ("head") causes the red-ward moving absorption while the high-velocity part ("tail") causes the blue-ward moving absorption simultaneously. The CIR model was worked out in detail by hydrodynamical computations by Cranmer & Owocki (1996). Fullerton et al. (1997) made another kinematic approach for
HD 64760, which agreed even better with the hydrodynamical model. In order to attempt to understand especially the variability in Hα we made a semi-empirical model in which all kinds of combinations of corotating structures can be included. A detailed description of this model is given in De Jong et al. (2000c, Chapter 7). We apply this model here to simulate the DAC behavior in the UV resonance lines of ξ Per.

Before we can apply our model, we must know the rotation period of ξ Per. It seems most obvious to take for this the most prominent period of 2.09 d. However, it turns out that subsequent DACs are not identical over this time interval. Especially the alternating column density (see Sect. 2.5.3 and Fig. 2.9, panel (c)) is a strong indication that the actual rotation period is twice as long, namely 4.18 d. De Jong et al. (1999a, Chapter 4) show that this is in agreement with the stellar parameters derived by Puls et al. (1996). With $v_{\text{rot}} \geq 204 \, \text{km/s} = v \sin i$, Penny (1996) the radius must be at least $17R_\odot$ and only their model with $M = 60M_\odot$ and $R = 25R_\odot$ complies with this. Leitherer (1988) derives a smaller radius of $12R_\odot$, which we cannot exclude at this stage. If we adopt $P_{\text{rot}} = 4.18$ d, this means that the corotating structure of ξ Per must contain at least 2 strong CIRs and 2 weaker ones in order to explain the strong and weak DACs observed within one cycle. We have computed such a model from which the structure is shown in Fig. 2.17 and the dynamic spectrum in Fig. 2.18. This dynamic spectrum resembles the observations quite well. The DACs start as weak and broad components, accelerate and become deeper and sharper. The banana-shaped isoflux contours are also present. However, our model does not include any perturbations to the velocity field. Cranmer and Owoczi (1996) found that the largest absorption takes place where the gradient in the velocity is lowest in a so-called “kink” in front of the CIR. This would mean that the maximum absorption from their model will occur earlier in time compared to our model.

2.7. Discussion

2.7.1. The CIR model

Nearly all observed wind variability down to the surface of the star can be explained by the presence of CIRs. The physical nature of the surface structure causing the CIRs is however by no means clear. In this observing run we have only the Hα and He I λ6678 lines that provide information on what happens near the stellar surface. The variations in these line profiles do not fit with a simple CIR model. The 2.09 d period in Hα is associated with the absorption features only between $-440$ and $-200 \, \text{km/s}$, i.e. in the wind. The periodogram of this line reveals no obvious periodic variations at higher velocities, but contains a collection of a number of weak maxima, which cannot easily be linked to regular variability (Fig. 2.12). The same can be said of the He I line (Fig. 2.13).
The Hα line also behaves quite differently compared to other years. During the MUSICOS campaign in November 1996 (Henrichs et al. 1998b, Chapter 3) there are no clear absorption events in the blue wing, but only red-ward moving absorption and emission events around the center of the line. Similar events were observed in the He I λ5876 line. The period analysis indicates that the longest timescale of variability is comparable to the length of the campaign. The highest peak does not belong to the 2 d period. In other Hα observations made in October 1991 a period of 2.0(1) d has been found between −200 and +50 km s$^{-1}$ where most of the variations occurred (Kaper et al. 1997). In observations of December 1998 a period of 2.06(4) d appears predominantly in the red wing (de Jong et al. 2000b, Chapter 5). The absorption in the blue wing was apparently much stronger in October 1994 than in other years in which ξ Per was observed.

We conclude that the timescale of the non-periodic variability in the optical lines is shorter than what is observed in the UV lines. This is probably because the formation of Hα is much more complicated than of the UV lines. The effects of small changes in the wind structure may have relatively a much larger impact on the Hα line-profile variations due to projection effects, cancellation of absorption and emission and the relatively small region in which Hα is formed. In ξ Per the Hα line is probably formed only up to 1.4R$_*$ (de Jong et al. 2000c, Chapter 7). Stochastic variations due to clumping may also play a role. In view of this it is interesting to note that the minimum of the edge EW at MBJD=8 coincides well with the maximum EW of Hα in the $[-500,-200]$ km s$^{-1}$ interval. In ξ Per this holds only for one event, but we note that in the O6I star Λ Cep a correlation between the He II λ4686 (formed close to the star) and C IV edge (far out in the wind) was observed in Oct 1989 during the whole observation run (Henrichs et al. 1990). According to the CIR model there should be a considerable phase lag between these regions. The correlation may be explained by assuming that the instability can accelerate material close to the star up to velocities which are comparable with the terminal velocity, thereby causing variability in the edge of the UV lines.

2.7.2. Non-radial pulsations and surface magnetic fields

We have investigated the presence of non-radial pulsations (NRP) and/or magnetic fields in ξ Per, which are the two most likely causes for azimuthal asymmetry at the stellar surface.

A NRP with a period of 3.45(2) h and mode $\ell=3$ was detected in a set of high time resolution spectra of the He I λ4713 line taken in October 1989 (de Jong et al. 1999a, Chapter 4). Some of the red-ward moving features in the He I λ6678 line of Oct 1994 and in the He I λ5875 line of Nov 1996 (results of the MUSICOS campaign, de Jong et al. 2000b, Chapter 5) are both very likely due to this phenomenon. To illustrate this we plotted in Fig. 2.19 a short timeseries of these lines and overplot the NRP phase maxima for comparison. The close correspondence means that this NRP mode is stable for at least 7 years. This particular NRP mode alone cannot be the cause of the observed cyclic wind variability, since the pattern speed is too high to be compatible with the wind period of 2.09 d. With $P_{\text{rot}} = 4.18$ d the corotating pulsation frequency is 6.2 c/d. If a "bright spot" with a CIR is associated with a certain pulsation phase, we would expect 6.2 CIRs per day crossing the line of sight for an equatorial observer. This frequency is too high to be compatible with the wind period of 2.1 days, although we cannot exclude that beating of several (not yet detected) pulsation modes may provide the low-frequency wind modulation. We think that the latter case is unlikely, however, because the main DAC frequency of all investigated O stars scales with $v_{\text{sin} i}$, whereas the (unknown) beating frequencies will not bear such a relation.

The magnetic field measurements presented here are not conclusive. The upper limit of 70 G in the disc-averaged longitudinal component of the field cannot exclude the presence of a field that is strong enough to perturb the flow. If we assume a dipole field which is tilted 90° with respect
to the rotational axis which has an inclination angle of 40° to the line of sight (de Jong et al. 1999a, Chapter 4) then the polar field strength ($B_p$) is about 6 times as large as the measured disk averaged longitudinal component of the field (de Jong et al. 2000b, Chapter 5). For a definite detection of a field the strength must be larger than 3r, which means that $B_p$ is larger than 1200 G. Such a field is more than strong enough to cause the CIRs. Magnetic fields are therefore still the most likely candidate for the wind variability, although such fields remain to be detected in O stars. For other attempts for $\zeta$ Per in November 1996 and December 1998 see Henrichs et al. (1998b, Chapter 3) and de Jong et al. (2000b, Chapter 5). For these attempts we have used the most sensitive instrument available today and obtained only upper limits. More sensitive instrumentation or techniques are apparently needed. We mention that in the B1 IIm star $\beta$ Cephei (Henrichs et al. 2000, Chapter 6) we recently found a sinusoidal varying magnetic field with a semi-amplitude of 90 G, which dominates the wind modulation, although differently than in O stars. More research in this direction is clearly needed before the enigma of the cyclical wind variability in early-type stars can be considered to be solved.

Acknowledgements. We thank the support received at the many observatories involved in this large project. We thank K. Annuk, L. Leedjärv and M. Ruusalepp for their help to observe at Tartu Observatory and S. Prins for her help to observe with the JKT. We thank Rens Waters for his helpful comments. JDJ acknowledges support from the Netherlands Foundation for Research in Astronomy (NFRA) with financial aid from the Netherlands Organization for Scientific Research (NWO) under project 781-71-053.

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Chapter 3

First results of the November 1996 MUSICOS Campaign on the O7.5III star \( \xi \) Persei

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Appeared in Proc. Cyclical variability in stellar winds, ESO workshop October 1997

Abstract. We present the first results of the MUSICOS campaign on the O7.5III star \( \xi \) Persei, held in November 1996, which was aimed to study its wind variability, rotation, pulsation and magnetic field in order to study their mutual effects. During 10 days at 8 observatories around the globe we obtained more than 300 high-resolution optical spectra between 4100 and 8000 Å, as well as magnetic field measurements from Hawaii and La Palma. So far we analysed the spectral lines of Hα, He I \( \lambda 5875 \) and O III \( \lambda 5592 \). CLEANed Fourier transforms of the three studied lines yield a complicated multiperiod behaviour and indicate that the most likely rotation period is about 4 days. Combining these data with data from earlier campaigns, we find strong evidence in the photospheric lines for prograde non-radial pulsations with a period of 3.5 h. Since the pulsation period is much shorter than the dominant cyclic period in the stellar wind features (as found in the UV lines, recorded in an earlier campaign including the IUE satellite), we can conclude that pulsation is very unlikely the driving agent for the cyclic wind variations, at least for \( \xi \) Per. The analysis of the magnetic field measurements is still in progress. Whether magnetic fields are responsible for the observed wind modulation can therefore not be answered at the present stage, but remains still the most likely option.

3.1. Introduction

All O and many B stars show systematic variability in their stellar winds, mainly in the absorption part of the UV resonance lines of Si IV, C IV, N V and other wind lines like N IV \( \lambda 1718 \). The most prominent features of variability are the discrete absorption components (DACs), which are migrating from low to high velocity towards the observer, with a recurrence timescale that can be interpreted as (an integer fraction of) the stellar rotation period.

The main unsolved issue is where the modulation comes from. A comparison of 4 timeseries of the Si IV profile of this star in subsequent years (Kaper et al. 1998) shows that the variations are cyclical with a dominant period of 2 or 4 days, but the variability slightly differs from year to year. Such behaviour strongly suggests a corotating pattern in the wind, but the pattern itself apparently changes on a timescale of less than a year. Other stars show similar long-term behaviour. See examples in Kaper et al. (1996 and 1999a).

The wind structures can be traced back to very low velocities: basically down to the vsini value of the star (see for \( \xi \) Per Henrichs et al., 1994 or Kaper et al. 1996). This argues in favour of a model with corotating windstructures, originating at or near the stellar surface, similar to Coro-
tating Interacting Regions in the Sun (CIRs, first proposed by Mullan, 1984). In the radiative hydrodynamical computations by Cranmer & Owocki (1996) these spiral-like regions indeed emerged, giving rise to accelerating DACs in the spectral lines, very similar to what is observed. They did not need to specify the origin of the perturbation: either magnetic fields or non-radial pulsations (NRP) could equally provide the required perturbation.

In order to find the origin of the observed wind variability we organized several campaigns on the O7.5III star ξ Persei, including magnetic field measurements. We choose this star because of its brightness and the very prominent DACs in the UV resonance lines with a suitable recurrence period of a few days. The timescales of pulsation, rotation and wind flow are all in the order of one day, which makes it particularly difficult to disentangle these effects, and a global network of observatories is needed to collect the necessary data. The strategy was to probe simultaneously the outer part of the stellar wind (UV resonance lines), the inner part of the wind (UV, N IV 1718, and optical Hα) and the stellar photosphere (optical O, Si and He lines). The deep photospheric lines are used to study pulsation behaviour by means of Doppler imaging techniques. The two most recent campaigns were in 1994, which included IUE and ground-based spectroscopy, and in 1996 (MUSICOS), with ground-based spectroscopy only.

The results of the October 1994 campaign are presented by de Jong et al. (2000a, Chapter 2). Here we describe the first results of the MUSICOS Campaign of November 1996, which was after the IUE satellite was taken out of service. The major improvement with respect to the previous campaign was the use of echelle spectrographs covering 4100 to 8000 Å, with some 20 suitable spectral lines, among which a number of deep photospheric lines which were not covered in the 1994 campaign, which included only the Hα and He I 6678 lines. These two lines appeared seriously wind contaminated, and therefore not suitable for pulsation studies.

With the MUSICOS spectropolarimeter at CFHT (Hawaii) and the SOFIN spectrograph at the NOT (La Palma) we also obtained surface magnetic field measurements. The instrumentation was about 5 times more sensitive than those used in October 1994 (Henrichs et al. 1998a, de Jong et al. 2000a, Chapter 2). The analysis of the magnetic data is still in progress. Here we present the first results from 3 optical lines, chosen to span a range of conditions in and above the photosphere.

3.2. The Campaign

We observed ξ Per during 10 days from November 17 till 27, with continuous coverage over 6 days and less coverage over 4 days. High S/N (> 300) echelle spectra where obtained at seven observatories (see Table 3.1). In total we obtained 313 spectra, out of which 176 are presently reduced and analysed. We used the MIDAS echelle package with an improved background subtraction for the INT and Xing-long spectra. The spectra of CFHT and OHP were reduced with for those instruments dedicated software. We removed the telluric lines from Hα and He I. We extracted, normalized and rebinned the regions around Hα, He I λ5875 and O III λ5592 on a uniform grid of 15 km/s. This resolution is an optimized compromise between the various resolutions to yield a sufficiently high signal to noise ratio (300–500) for the kind of variations we are looking for. Most of the spectra taken simultaneously at different observatories match within the achieved accuracy.

3.3. Results

The dynamic quotient spectra of Hα, He I λ5875 and O III λ5592 are shown in Fig. 3.1. The general parallel behaviour between the patterns in Hα, He I and O III is apparent, notably around \( t = 8 \) and \( t = 9.2 \) days. Note also the sharp, linear pattern around \( t = 12 \) in He I and O III (not in Hα), which appears only between −200 and 200 km/s (\( = v \sin i \)) and migrates very rapidly from blue to red.

**Period analysis**

We performed a CLEANed Fourier Transform period analysis on the 3 lines. Fig. 3.2 shows the powerdiagrams of Hα and He I. At this stage of the analysis we do not trust the period analysis for the O III line because of remaining systematic differences between simultaneous spectra from different observatories. In Hα the most prominent frequencies appear at 0.18(9) c/d (5.5 d), 0.44(6) c/d (2.3 d) and 1.42(6) c/d (0.70 d). The large uncertainty in the first period is due to the relatively short coverage. For He I we find similar periods: 0.20(6) c/d (4.4 d), 0.49(10) c/d (2.0 d), 1.02(11) c/d (0.98 d) and 1.47(12) c/d (0.68 d). The given uncertainties are calculated from the widths of the peaks in the powerdiagrams. This has to be compared with the period analysis of our previous campaign in October 1994, where we found periods of 2.08 d, 1.04 d and 0.67 d in the Hα, Si IV and N IV lines, respectively (Henrichs et al. 1998a, de Jong et al. 2000a, Chapter 2).

**Non-radial pulsation**

We have investigated the possibility that the fast moving absorption feature around \( t = 12.2 \) in the weak photospheric lines O III λ5592 and He I λ5875 is due to NRP. The steepness indicates a few-hour period. Such a period is however too short to determine with certainty from this campaign because of the sampling rate. After reanalysing previous observations of the weak photospheric He I λ4713 line taken in October 1989 (from Calar Alto and Kitt Peak) and joining these with the MUSICOS results, we found a very clear NRP period of 3.5h, \( l = 3 \) (or less likely \( l = 4 \)) mode and 0.2% full amplitude (de Jong et al., 1999a, Chapter 4).

**Magnetic fields**

The bad wheather conditions at Hawaii prevented that
Table 3.1. Overview MUSICOS November 1996 Campaign on ξ Per

<table>
<thead>
<tr>
<th>Observatory</th>
<th>Location</th>
<th>Observers</th>
<th>Coverage</th>
<th>Spectra</th>
</tr>
</thead>
<tbody>
<tr>
<td>CFHT</td>
<td>Hawaii, USA</td>
<td>Boehm, Catala, Donati Landstreet</td>
<td>4101–8138Å</td>
<td>36</td>
</tr>
<tr>
<td>INT</td>
<td>La Palma, Canary Islands</td>
<td>de Jong, Ehrenfreund, Foing, Oliveira, Stempels, Teltin</td>
<td>4305–9490</td>
<td>50</td>
</tr>
<tr>
<td>McDonald</td>
<td>Texas, USA</td>
<td>Hatzes, Johns-Krull, Neff</td>
<td>5430–6723</td>
<td>76</td>
</tr>
<tr>
<td>OHP</td>
<td>France</td>
<td>Henrichs, Schrijvers</td>
<td>3892–6817</td>
<td>42</td>
</tr>
<tr>
<td>Ritter Obs.</td>
<td>Toledo, USA</td>
<td>Mulliss</td>
<td>5440–6827</td>
<td>9</td>
</tr>
<tr>
<td>Xinglong</td>
<td>Beijing, China</td>
<td>Hao, Huang, Yang</td>
<td>5510–8389</td>
<td>53</td>
</tr>
<tr>
<td>Ondrejov</td>
<td>Czech Republic</td>
<td>Cao</td>
<td>6300–6700</td>
<td>6</td>
</tr>
<tr>
<td>NOT</td>
<td>La Palma</td>
<td>Dümmler, Ilyin</td>
<td>3580–10800</td>
<td>41</td>
</tr>
</tbody>
</table>

Fig. 3.1. The dynamic spectra of the lines Hα, HeI 5875Å, and OIII 5592Å. The upper panels show the average spectra. The middle panels show an overplot of the quotient spectra. In the greyscale images in the lower panels we can see that the features in the three different lines are quite similar and partly probe the same region. The horizontal axis is drawn in the stellar rest frame.

3.4. Discussion and conclusions

The dominant, well-known period of ~ 2 d of the UV wind lines is confirmed by the MUSICOS data in all studied (optical) lines, which therefore supports the hypothesis of a rotation modulation of the wind. It is interesting to note that the behaviour of the base of the wind looks similar, but not identical to what we observed in 1994 (see de Jong et al. 2000a, Chapter 2 or Henrichs et al. 1998a for a dynamical spectrum). The inner wind structure has apparently slightly changed its configuration.
Although the long periods in Hα (5.5 d) and He I (4.4 d) are very uncertain, they represent a real timescale of variability, since the observed variations cannot be reproduced without such modulations. Within the errors these are compatible with a 4 day rotation period which is twice the period we found earlier in the UV and Hα (Henrichs et al. 1998a, de Jong et al. 2000a, Chapter 2). This period would require a minimum radius of 16R_☉ which falls within the range as calculated from atmospheric model fits for this star (Puls et al. 1996). We therefore adopt the 4 day period as the rotation period.

Summarizing, the 3.5 h NRP period, as derived from earlier data, is too short to cause directly the 2 or 4 day wind modulation period. The combined effects of such pulsations and rotation on the wind need to be investigated.

The presence of small surface magnetic fields remains the most plausible hypothesis for modulating the stellar wind, but this is still to be demonstrated. A simple estimate yields that 50–100 G would be sufficient to perturb the wind.

Acknowledgements. We are very much indebted to the scheduling staffs of all observatories for their help to coordinate this campaign. Part of this research is supported by the Netherlands Organization for Scientific Research (NWO project nr. 781-71-053).

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Chapter 4

Non-radial pulsations in the O stars ξ Persei and λ Cephei

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Appeared in A&A 1999, volume 345, page 172

Abstract. A new time-series analysis of profile changes in the photospheric He I λ4713 spectral line from data taken during 5 days in 1989 at the Calar Alto and Kitt Peak observatories has provided evidence for the presence of a non-radial prograde p-mode in the O 7.5 giant ξ Per (ℓ = 3, P = 3.5) and probably two such modes in the O6 supergiant λ Cep (ℓ = 3, P = 12.3 h and ℓ = 5, P = 6.6 h). The rotating pulsation periods are in both cases much shorter than the estimated stellar rotation period. Modeling the observed amplitude of the line-profile changes (assuming |m| = ℓ) yields a velocity amplitude of approximately 5 km s⁻¹ of the pulsation in ξ Per and 6 km s⁻¹ in λ Cep.

Any such a pulsation by itself is unlikely to be the cause of the well-known cyclical wind variability in these stars because the pulsation period is too short, but the cumulative action of multiple modes could cause such an effect. Weak magnetic fields anchored at the surface remain the strongest candidate for the origin of wind variability.

4.1. Introduction

Pulsating stars are found in nearly every part of the HR diagram. Among O stars, however, only very few pulsators are known. Fullerton et al. (1996) listed 3 confirmed and 6 suspected pulsating O stars and noticed that all these massive stars are located in some cases in the neighborhood of the Corotating Interaction Regions (CIRs), where postulated coronal mass ejections can emerge. The CIR model invokes fast and slow wind streams that originate at different locations at the stellar surface. Due to the rotation of the star, the wind streams are curved, so that fast wind material catches up with slow material in front forming a shock at the interaction region. The shock pattern in the wind is determined by the boundary conditions at the base of the wind and corotates with the star.

The wind structures can be traced in radial velocity back down to the v sin i value of the star (see for ξ Per Henrichs et al. 1994, Kaper et al. 1996). This argues in favor of a model with corotating wind structures similar to the Corotating Interaction Regions in the solar wind (CIRs), in the context of hot-star winds first proposed by Mullan (1984). The CIR model invokes fast and slow wind streams that originate at different locations at the stellar surface. Due to the rotation of the star, the wind streams are curved, so that fast wind material catches up with slow material in front forming a shock at the interaction region. The shock pattern in the wind is determined by the boundary conditions at the base of the wind and corotates with the star.

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The timescales of pulsation, rotation and wind flow are all on the order of one day, which makes it particularly difficult to disentangle these effects, and which forces a ground-based multi-site approach, preferably simultaneously with UV spectroscopy from space. A different approach has been followed by Howarth et al. (1998), who were able to recover pulsation periods of HD 64760 B0.5 Ibn and HD 93521 O9.5V derived from UV data by applying cross-correlation techniques. Their paper is concerned with the same problem as addressed here. In our study we decided to concen-
tate on the O stars ξ Persei O7.5III(n)(f) and λ Cephei O6I(n)p because of their brightness, conveniently close relative location in the sky, excellent record of their UV resonance line behavior, and suitable recurrence period of DACs (1–2 days). The strategy was to probe simultaneously the outer part of the stellar wind (using UV resonance lines), the inner part of the wind (using N IV λ1718 and HeII) and the stellar photosphere (using optical lines). Lines formed deep in the photosphere are used to study pulsation behavior by means of Doppler imaging techniques.

This paper reports the results of a new analysis of the behavior of such a photospheric line He I λ4713 of such a campaign in 1989 during 5 days. Whereas previous analyses of this dataset did not yield convincing results, the advent of new methods of line-profile analysis allowed us to detect the pulsation modes described below.

**4.2. Observations and data reduction**

A multi-site campaign was held from 17 to 22 October 1989, including UV spectroscopy with the IUE satellite, optical spectroscopy at the Kitt Peak and Calar Alto observatories, B polarimetry at McDonald Observatory and UBV photometry at the Wendelstein observatory. The IUE spectra are described and analyzed by Kaper et al. (1996, 1999a). Here we restrict ourselves to the data that lead to the pulsation analysis.

The experimental setup was chosen in accordance with the vini values of the stars. With 200 km s⁻¹ for ξ Per and 214 km s⁻¹ for λ Cep (Conti & Ebbets 1977) a resolution of 15 km s⁻¹ would give about 25 effective points in the line profile, sufficient to assure that modes up to ℓ ≈ 20 would be detectable. We concentrated on the weak and therefore deep-seated photospheric He I absorption line at 4713 Å.

The Calar Alto 2.2m telescope (observers GZ and HH) was used with the 90cm f/3 camera in Coude focal with gratings #1 (632 lines mm⁻¹) in second order centered at 4667 Å, giving a dispersion of 8.6 Å mm⁻¹. We used a RCA CCD chip (#11) with 1024×640 pixels of 15μm size yielding a coverage of 130 Å with a 2 pixel resolution of 0.26 Å. At this setting the CCD has a quantum efficiency of 68%, which gave 6.7 electrons per count with the gain set at 30. The slitwidth was 1.5 arcsec, equal to the average seeing. At Kitt Peak (observer DG) the 0.9-m Coude feed was used with camera #5, long collimator with grating A (632 lines mm⁻¹) in second order with a 4-96 filter to block the first order light, yielding a dispersion of 7.1 Å mm⁻¹ on the Texas Instruments CCD (T13) with 800×600 pixels of 15μm size. This yielded a coverage of 85 Å with a projected slitwidth of 21.3μ FWHM. The CCD chip has a quantum efficiency of 70%; giving 4.12 electrons per count with the used gain. The slitwidth was fixed at the average seeing of 1.5 arcsec. Typical exposure times were 2 to 3 minutes for ξ Per and 7 to 10 minutes for λ Cep at Calar Alto, and about 1.5 times longer at Kitt Peak. The number of spectra used in the analysis was 324 for ξ Per (154 from Calar Alto and 170 from Kitt Peak), and 169 for λ Cep (109 from Calar Alto and 60 from Kitt Peak). Fig. 4.1 depicts the time line of the optical spectroscopic observations for both stars.

Flatfielding, bias subtraction and wavelength calibration with Th–Ar lamps were done in the usual manner. The spectra were normalized by using a linear fit through wavelength segments carefully selected to contain no traces of spectral lines. We used [4695.2, 4700.4] and [4720.1, 4723.4]Å for both stars. We also attempted higher-order polynomials, but this did not change the results of the frequency analysis. All spectra were finally smoothed and reduced on a uniform velocity grid of 15 km s⁻¹. Spectra taken at the same epoch at the two observatories showed excellent agreement within the errors given the signal to noise ratio (S/N) which was on average about 800 (Calar Alto) and 500 (Kitt Peak) for ξ Per and 500 (Calar Alto) and 300 (Kitt Peak) for λ Cep, respectively. This agreement was reached after some small night-to-night and observatory-to-observatory wavelength corrections. These were only needed for Calar Alto spectra, because of some technical problems at the telescope that occurred at several nights, which made apparently small changes to the wavelength setting.

Throughout this paper, the reference wavelength for conversion to the stellar restframe has been corrected with 60 and −75 km s⁻¹ for the runaway stars ξ Per and λ Cep, respectively (Gies 1987).

Average spectra are shown in the top panels of Figs. 4.2 and 4.3. Sample quotient dynamic spectra of ξ Per and λ Cep are in these figures compared with the inverse Fourier transform (after removal of noise) and with folded data using the pulsation frequencies reported in this paper.

**4.3. Period analysis**

We performed a standard CLEAN analysis (Roberts et al. 1987, Gies & Kullavanijaya 1988) of the He I λ4713 Å line. The resulting power spectra are shown as grey-scale representations in Figs. 4.4 and 4.6. There is considerable power near periods around 1 day and longer, which for both stars we can mostly attribute to wind contaminations, rather than windowing aliases. These low-frequency signals may
Non-radial pulsations in the O stars \( \xi \) Persei and \( \lambda \) Cephei

Fig. 4.2. Figure at the left: dynamic quotient spectra of 2 nights of \( \xi \) Per data. Horizontal lines indicate the mid-exposure times. The upper panel shows the average spectrum used to create the quotient spectra which are shown in the panel below. Note the slight depression at the left part of the average spectrum, caused by an unidentified line. Middle figure: inverse Fourier transform for which frequencies only above 2 cd^{-1} were used (see Sect. 4.3). Figure at the right: all data folded with a 3.5 h period, rebinned to 25 points per cycle

also partly reflect unaccounted systematic differences between observatories and slight differences in processing of flat fields, normalization, etc. between nights, in spite of our efforts to remove them. In the following we ignore this frequency range. In the uncontaminated areas (above 1.5 cd^{-1}) significant power is found across the line profile at several frequencies, visible as horizontal grey areas in the figures, which we attribute to NRP. The frequencies quoted in Sect. 4 were determined by first fitting in each velocity bin a Gaussian function to the peak in the power \( P(f) \):

\[
P(f) = P_{\text{max}} \exp(- (f - f_0)^2 / 2 \sigma_f^2).
\]  

(4.1)

This yields three parameters: the maximum power \( P_{\text{max}} \), the central frequency \( f_0 \) and the width of the peak \( \sigma_f \). To determine the uncertainty \( \epsilon_f \) in \( f_0 \) the height of the peak relative to the noise level, defined as \( N^2 \), should be taken into account. For this purpose we use the width of the peak at \( P_{\text{max}} = N^2 \), following Schwarzenberg-Czerny (1996). Applying Eq. 4.1 gives:

\[
\epsilon_f = \sigma_f \sqrt{-2 \ln \left(1 - \frac{N^2}{P_{\text{max}}} \right)}.
\]

(4.2)

We use a conservative method to calculate the noise level, which is a critical parameter: we defined \( N^2 \) as the power below which 95% of the datapoints fall in the histogram of the periodogram between 5 and 10 cd^{-1}. For \( \xi \) Per we find \( N^2 = 2.4 \times 10^{-8} \) and for \( \lambda \) Cep \( N^2 = 2.2 \times 10^{-8} \). We have chosen this procedure because white noise gives low and high peaks at all frequencies and apart from this it is not always clear whether a peak is only noise or contains a weak periodic signal.

Finally, the average frequency across the line profile was computed in two ways: a normal average with its standard deviation and an error-weighted average. In the latter
method the uncertainty in the average is directly derived from the errors, $\epsilon_f$, using the following equations:

$$\bar{f} = \frac{\sum_{i=1}^{N} f_i / \epsilon_f}{\sum_{i=1}^{N} \epsilon_f^{-2}},$$

$$\bar{\epsilon}_f^2 = \frac{1}{\sum_{i=1}^{N} \epsilon_f^{-2}}.$$  \hspace{1cm} (4.3)

(4.4)

In both cases we computed for every datapoint the deviation $d_i = |(f_i - \bar{f})/\sigma_f|$ where $\sigma_f$ is the standard deviation. Secondly, we computed for each point the probability $P(d_i)$ of being a statistical fluctuation assuming a Gaussian distribution. We discarded all points for which $N \cdot P(d_i)$ is smaller than 0.5 (Chauvenet’s criterion). In a sound statistical distribution both methods should give the same values. In case of differences we have chosen the method giving the largest error.

The phase of the signal can be extracted from the CLEANed Discrete Fourier Transform (CDFT), using the following representation:

$$F(t) = \Delta f \sum_{i=1}^{N} \sqrt{2P_i} \cos(2\pi[f_i(t - \bar{t}) + \phi_i]).$$  \hspace{1cm} (4.5)

Here $F(t)$ is the flux as a function of time, $t$, $f_i$ is the frequency, $\Delta f$ is the frequency interval $f_{i+1} - f_i$, $P_i$ is the power, $\phi_i$ is the phase of component $i$, and $\bar{t}$ is the average time of the sample, used to relate the calculated phases to the Barycentric Julian Date (2447818.698 for $\xi$ Per and 2447818.532 for $\lambda$ Cep). By convention $\phi$ is defined between 0 and 1. Note that features with larger phase arrive earlier in time.

As an independent check we derived the phase information also from least-$\chi^2$ multiple sine fits (components in Eq. 4.5) with frequencies fixed on the main peaks in the periodograms. We used 0.997, 9.138 and 7.956 c d$^{-1}$ for $\xi$ Per and 0.424, 1.96, 2.54 and 3.67 c d$^{-1}$ for $\lambda$ Cep. Error bars on the phases could then be derived by means of a Monte Carlo method in which artificial noise was added within the S/N limits of the data. In Figs. 4.5 and 4.7 we show the standard deviations $\sigma_\phi$ and $\sigma_A$ of the phases and amplitudes resulting from 25 fits for each point. Before computing $\sigma_\phi$ the phases ($\phi_i$) were anchored to the first calculated value using $\phi'_i = \phi_i - \text{INT}(\phi_i - \phi_0 + 0.5)$. This assures that any value of $\phi'_i$ differs less than 0.5 from $\phi_0$. The error bars are therefore not expected to exceed $\sqrt{1/12} \approx 0.28$ which is the standard deviation of a uniform distribution with phases between 0 and 1 in absence of any signal. The apparently non-random distribution of the values of some phases outside the line profiles in Figs. 4.5 and 4.7 is unclear. We think that this might be caused by very small residual periodic variations in the continuum as seen in Figs. 4.2 and 4.3, rather than by a statistical dependency of the different points.

Fig. 4.3. Figure at the left: dynamic quotient spectra of 3 nights of $\lambda$ Cep data, in similar format as in Fig 4.2. Middle figure: inverse Fourier transform using frequencies only between 1 c d$^{-1}$ and 5 c d$^{-1}$. Figure at the right: sum of data folded by respectively 12.3 and 6.6 h and rebinned to timesteps of 0.0075
The quoted amplitudes and phases are determined from these multiple sine fits. These amplitudes are more reliable than from the CDFT's. This is because CLEAN removes the window-function effects component by component from the DFT, regardless whether the peak being removed contains some power of a real frequency. Consequently amplitudes in a CDFT may be smaller than they truly are. The phases determined by the two methods agree within the error bars.

As a comparison we reconstructed the pulsation patterns by inverting the DFTs and added all the components of Eq. 4.5 within the frequency range of interest, thereby filtering out a large fraction of the noise. We also folded the data with the detected periods. These two resulting reconstructions, in a grey-scale representation and appropriately rebinned, are shown in Figs. 4.2 and 4.3 along with the original data.

As an independent method for finding periods, complementary to the CLEAN analysis, we also performed a minimum-entropy method (Cincotta et al. 1995). In this method the time series are folded with trial periods, i.e. the phase of each observation is computed as \( \phi = t / P - \int \text{INT}(t / P) \). The phase/flux space is then divided into a number of grid elements in which the number of datapoints is counted. By defining the probability of finding a data-point in element \( i \) as \( p_i = N_i / N \) the entropy is calculated as \( S = - \sum p_i \ln p_i \). The lower the value of \( S \), the higher the probability that the trial folding period corresponds to a true period. The significance of the considered period follows from \( \sigma^2 = (S - S_0)^2 / \sigma_c^2 \), where the index \( c \) refers to the continuum. In order to deal with the window function caused by the inhomogeneous data sampling, the minimum-entropy method was first performed after shuffling the data in a random manner, yielding \( S_0 \). Subtraction of \( S_0 \) from \( S \) removed in this way the periods caused by the window function.

4.4. Application and results

4.4.1. \( \xi \) Per (HD 24912)

In the periodogram (see Fig. 4.4) the most significant power outside the contaminated frequency range is found at \( f_1 = 6.96(4) \, \text{c d}^{-1} \) (or \( P_1 = 3.45(2) \, \text{h} \)) and at \( f_2 = 9.14(7) \, \text{c d}^{-1} \) (or \( P_2 = 2.63(2) \) h). The corresponding amplitudes and phases are displayed in Fig. 4.5. The phases for amplitudes less than \( 2 \sigma \) above the continuum are not significant and are plotted as open symbols. We argue below why we do not consider \( f_2 \) as a significant period, and first consider \( f_1 \). Whereas the phase behavior of \( f_1 \) is clearly characteristic for a NRP, the asymmetry of the amplitude around line center is not, unless non-adiabatic temperature variations are important. In the profiles of \( \xi \) Per we find small equivalent-width (EW) changes, which could be in principle due to temperature effects. A CLEAN analysis of the EW variations show a few weak peaks in the power diagram, none of which correspond to any other known frequency in this star. The source of the EW variability is therefore probably noise. Temperature effects generate in general larger amplitudes in EW than we observe (although cancellations cannot be excluded), which makes us conclude that these effects are not likely to play a significant role (see Schrijvers & Telting 1998). Data with a higher S/N and better time resolution are needed to establish the (expected) first harmonic and to determine the anisothermal number \( m \).

We investigated whether the relatively low power on the blue side at \( f_1 \) could be due to wind contamination. Simultaneously taken UV spectra (Kaper et al. 1999a) show that a new DAC developed at low velocity around BJD 2447818.9. Similar simultaneous observations from another campaign on this star (see Henrichs et al. 1998a) showed that the new development of a DAC is accompanied by enhanced blue-shifted absorption in H\( \alpha \). The He I line presently studied is probably also partly formed in

![Fig. 4.4. Grayscale representation of the periodograms as a function of velocity of \( \xi \) Per spectra. The side panel shows the power summed over all velocities. The highest peak is normalized to unity. The top panel displays the ratio of the observed standard deviation to the expected standard deviation.](image-url)
the wind and therefore should in principle display similar kind of extra absorption, which would disturb the period analysis. We therefore excluded the 24 spectra between BJD 2447818.975 and BJD 2447819.1 in our further period analysis. These spectra showed extra blue-shifted absorption which was not present in other spectra, and which we attribute to wind effects. This procedure indeed decreased somewhat the asymmetry in amplitude, although not completely. Some wind absorption is undoubtedly still present in a number of spectra, but a thorough elimination is beyond hope. A second cause might be a blending effect by the partly overlapping (unidentified) weak lines in the alleged blue continuum which may have distorted the normalization. We consider this as less important.

We furthermore folded the spectra with $P_1$ (right panel in Fig. 4.2). This shows a clear NRP pattern in the range from $-100$ to $160$ km s$^{-1}$, less asymmetric around zero velocity than Figs. 4.4 and 4.5 suggest. Although the asymmetry in power remains worrisome, more evidence for $P_1$ being a true period in this star comes from observations of the O III A5592 line taken during the MUSICOS November 1996 campaign (see Henrichs et al. 1998b, Chapter 3), which also clearly show the NRP pattern over the whole line profile. This dataset was however insufficiently sampled to determine any NRP frequencies.

Considerable power is also found at $f_2 = 9.14(7)$ c d$^{-1}$ (or $P_2 = 2.63(2)$ h) over almost the whole line profile (see Fig. 4.4). Several reasons argue against an interpretation in terms of NRP, however. First, the minimum-entropy analysis (see Sect. 4.3) did not reveal any signal around $f_2$, whereas $f_1$ was detected with $15\sigma$ significance. Secondly, the phase behaviour is suspiciously similar to that of $f_1$ (see Fig. 4.5). In addition, the amplitude increases where the amplitude of $f_1$ decreases, which may point towards interference. For these reasons we do not consider $f_2$ as a NRP frequency.

From the phase diagram belonging to $f_1$ we can derive the azimuthal degree, $\ell$, according to Telting & Schrijvers (1997, hereafter TS) by measuring the difference in phase at $\pm 230$ km s$^{-1}$, i.e. just outside $\pm v\sin i$ of the star. We note that our adopted $v\sin i$ value is in accordance with Penny (1996), although Howarth et al. (1997) find $213$ km s$^{-1}$. From Fig. 4.5 we derive $\Delta \Psi (f_1) = 2\pi \times 1.6(2)$, where we had to use a linear extrapolation of a fit through the region between $-125$ and $180$ km s$^{-1}$ because outside these regions the phase is likely to be poorly defined due to the very low amplitudes.

Table 2 of TS lists coefficients of empirical linear fits to input and output $\ell$ values of synthetic data generated by Monte Carlo calculations for various pulsation parameters. We show below that in our case the ratio of horizontal to vertical motions, $k$, is small, certainly less than 0.3. This constraint gives for $\Delta \Psi (f_1)$ the result $\ell_1 = 3.4(5)$ with $88\%$ confidence that the $\ell$ value is correct within $\pm 1$. We adopt $\ell_1 = 3$ as the most probable value for the harmonic degree. A higher or lower $\ell$ value would show up as one more or one less absorption feature in the line profile, which is not seen (Fig. 4.2). Since the first harmonic of this signal (at 14 c d$^{-1}$ or 1.75 h) could not be detected in our dataset because the sampling rate was too low, we have no means to derive a value for the azimuthal parameter $m$.

We can determine the direction of the pulsation mode if the stellar rotation rate is known. Periodicities in stellar wind features suggest that the rotation period, $P_{\text{rot}}$, is 2 or 4 days (Kaper et al. 1999a). We obtain in both cases that the mode must be prograde. The value of $k$ is in the (here justified) Cowling approximation related to the pulsation frequency, $f_{\text{co}}$, in the corotating frame: $k = GM/R^2(2\pi f_{\text{co}})^2$. Two rather differing values for the mass and radius are known: Leitherer (1988) finds 35 M$_\odot$ and 12 R$_\odot$ (case 1), not differing very much from earlier determinations, and also in agreement with the values given by Howarth & and Prinja (1989), whereas Puls et al. (1996) derive 60 M$_\odot$ and 25.5 R$_\odot$ (case 2) with new model-atmosphere fits, including wind effects. Even with the distance limits set by HIPPARCOS no preference can be given to either set. We can exclude a 4-day rotation period in case 1 because the implied rotation velocity would be lower than the observed $v\sin i$ value, whereas a 2-day period is excluded in case 2, since the star would rotate at break-up. In both remaining combinations (case 1 with $P_{\text{rot}} = 2$ d and case 2 with $P_{\text{rot}} = 4$ d) the inclination angle turns out to be close to 40°. The corresponding corotating frequencies, derived
Non-radial pulsations in the O stars ξ Persei and λ Cephei

4.4.2. λ Cep (HD 210839)

Outside the contaminated frequency range we find significant power at two different frequencies: at \( f_1 = 1.96(8) \) c d\(^{-1}\) (or \( P_1 = 12.3(5) \) h) and at about twice this frequency at \( f_2 = 3.64(14) \) c d\(^{-1}\) (or \( P_2 = 6.6(3) \) h), see Fig. 4.6. The minimum-entropy analysis yielded for these signals a significance of 10\( \sigma \) and 8\( \sigma \), respectively. The corresponding amplitudes and phases are displayed in Fig. 4.7. The power of the first period appears to be concentrated mostly on the blue side of the line. This could be caused by interference of the unidentified line which is displaced by about -400 km s\(^{-1}\) with respect to the He I line, and which is stronger than in ξ Per. This secondary NRP pattern can clearly be seen in the folded spectra (Fig. 4.3). From the phase diagram of the larger period we derive an \( \ell \) value with the same method as above by measuring the difference in phase at ±250 km s\(^{-1}\), again just outside ±vsini of the star. Our adopted vsini value is in accordance with Penny (1996), not very different from 217 km s\(^{-1}\) given by Howarth et al. (1997). We find \( \Delta \Psi(f_1) = 2\pi \times 1.3(1) \). The irregular phase jumps near -160 km s\(^{-1}\) are likely caused by the very low power at these velocities, which are probably due to imperfection of our dataset. We use here the \( \ell - |m| < 2 \) restricted fit, which gives \( \ell_1 = 2.9(3) \) according to TS. We adopt therefore \( \ell_1 = 3 \) as the most probable value.

Although \( f_2 \) is nearly twice \( f_1 \), it cannot be its first harmonic since the ratio \( f_2/f_1 = 1.86(10) \) deviates more than 1\( \sigma \) from the exact value of 2. In addition, in all velocity bins the power at \( f_2 \) lies systematically below the power at \( 2f_1 \) (see Fig. 4.6). This makes the probability of \( f_2 \) being a harmonic of \( f_1 \) less than 10\(^{-5}\). We could also exclude \( f_2 \) being a harmonic by considering the consistency check for the amplitude ratio and the phase relation of the main frequency and its first harmonic as given by TS. They find that \( \Psi_{12} = 2\Psi_1 - \Psi_2 = \pi \times 1.50(6) \), where the phase at line center of the main frequency is denoted by \( \Psi_1 \) and of the first harmonic by \( \Psi_2 \). For Λ Cep we obtain \( \Psi_{12} = \pi \times 1.6(3) \). This would give fair confidence that the higher frequency could indeed be the first harmonic of the lower, except that in all model calculations by TS the ratio of the amplitudes of the first harmonic to the main frequency is considerably...
smaller than we observe, which makes $f_2$ as a first harmonic very unlikely, which we therefore consider as a second NRP mode.

For the second mode we find $\Delta \Psi(f_2) = 2\pi \times 2.4(3)$. From this value we derive $i_2 = 5.2(7)$ using a $k < 0.3$ fit with 86% confidence, which implies $\ell = 5$ as the most probable value for this second NRP mode and $m$ remains undetermined. This mode complies with the number of bumps seen at any given time in the folded spectra in Fig. 4.3.

Adopting an upper limit to the rotation period of 4.5 days, using a radius of 19 R$_\odot$, a mass of 59 M$_\odot$ (Puls et al. 1996) and an inclination angle of 90°, we find that the corotating frequencies are 1.29 cd$^{-1}$ for the $\ell = 3$ mode, implying $k < 0.38$, and 2.53 cd$^{-1}$ for the $\ell = 5$ mode, corresponding to $k < 0.1$. Both modes are therefore prograde modes, and could be a $p$ or $g$ mode, depending on the adopted stellar model.

### 4.5. Model calculations

From the observed amplitudes in the line profiles one can derive the velocity amplitude of the pulsation. To obtain a crude estimate we considered for simplicity a single sectoral $p$-mode ($\ell = |m|$), i.e. without horizontal velocity component ($k = 0$), and no temperature variations. We divided the visible stellar surface in about 2500 elements. The grid is further defined by the inclination angle, $i$, rotational velocity, $v_{rot}$, and pulsation velocity semi-amplitude $\Delta v$. For each grid element we calculated specific intensities and limb darkening of the He I $\lambda 4713$ line for the given stellar parameters ($T_{\text{eff}}$ and $\log g$) using atmosphere models created by TLUSTY and SYNSPEC by Hubeny & Lanz (1992). For each grid element the radial velocity and angle towards the observer is computed. Using these values, Doppler-shifted specific intensities as function of $v$ are interpolated from the model atmosphere. Finally, the line profile was calculated by integrating the intrinsic line profiles of all grid elements over the visible surface. For a given inclination angle and pulsation amplitude 20 profiles were calculated spread over a full period and the profile with the largest amplitude at line center was selected. At that particular phase the inclination and $\Delta v$ were varied in the domain of Fig. 4.8 and the maximum amplitude was measured. We applied this method to $\xi$ Per with $T_{\text{eff}} = 36 000$ K and $\log g = 3.4$ (Puls et al. 1996, case 2 in the above). We computed 25 models in the $\Delta v - i$ plane between 5 and 25 km s$^{-1}$ and from 10 to 90° respectively. The resulting semi-amplitudes relative to the line depth, $A_{\text{rel}}$, are shown as contours in Fig. 4.8.

In $\xi$ Per the central depth of the He I line is 4% and the amplitude only 0.12% of the local continuum (see Fig. 4.5), which means that the NRP signal is weak with $A_{\text{rel}} = 0.03$. Following the corresponding contour in Fig. 4.8 we derive that $\Delta v$ can be at most about 5 km s$^{-1}$ for the adopted $i \sim 40°$.

For $\lambda$ Cep the stellar parameters are not too different from those of $\xi$ Per (although $k$ is larger), and we simply applied the same calculations for this star as a first approximation. The central depth of the line is 2% and the amplitude 0.11%, implying $A_{\text{rel}} = 0.06$. For an inclination angle of $i \sim 90°$ this gives $\Delta v \sim 6$ km s$^{-1}$ for this star. The inclination cannot be much lower according to the stellar parameters (see Sect. 4.4.2). From a sample calculation we found that a model atmosphere for $\lambda$ Cep has nearly the same limb darkening and width of the intrinsic profiles as for $\xi$ Per, which are the main quantities on which the NRP amplitude depends. In spite of all these approximations we consider the derived value for $\lambda$ Cep to be quantitatively justified.

### 4.6. Discussion

The length and coverage of our datasets, combined with known cyclical wind effects limit the frequency range in which we can detect pulsation periods from about 1.5 to 10 cd$^{-1}$. Of course we cannot rule out the presence of other modes in these stars. In fact it is likely that more modes will be found with better datasets like for example in the case of the 09.5V star $\zeta$ Oph (Kambe et al. 1993, Reid et al. 1993, Kambe et al. 1997, Jankov et al. 1998) where increasingly higher-quality spectra and denser coverage revealed a larger number of modes, up to $\ell = 18$.

The presence of NRP in $\xi$ Per was already suspected by Gies & Bolton (1986) on the basis of radial velocity variations, which they assumed to be due to NRP, but no significant period or mode could be identified from their sample of 38 photographic spectra. Significant profile variability in many lines (except in $\lambda 4713$ Å) was also found by Fullerton (1990), but no mode or period could be established.
A preliminary analysis of the data of the present paper on $\lambda$ Cep was given by Henrichs (1991), who found the $\ell = 5$ NR P mode with a 6.5 h period, and mentioned the possible presence of the $\ell = 3$ mode at 12.5 h. Walker (1991) reported a likely $\ell = 5$ or 6 mode (no period) for this star, but without further details we cannot compare his results with ours.

We also find small EW variations in this star, and some frequencies do coincide approximately with the pulsation frequencies. They could in principle be to be due to temperature effects. Higher-quality data are needed to confirm this.

Marchenko et al. (1998) report the presence of a strong 0.63 day period in the HIPPARCOS photometry of $\lambda$ Cep (but none in $\xi$ Per) over the 3.5 year lifetime of the satellite, which partly covered our campaign. This period is not compatible with any of the NR P or wind periods, and its origin is unclear.

We now return to the question set out at the beginning of this research whether non-radial pulsation can be the origin of the cyclical wind variability. If the mode we found is the only mode in $\xi$ Per, it appears that the pattern speed of the waves running around the star superposed on the rotation (see Sect. 4.4.1) is too high to be compatible with the observed wind periods (1, 2, 4 days). This is probably true for most short-period single mode pulsations, since for most O stars the rotation rate is much lower. An exception is the O4 star $\zeta$ Pup, for which Reid and Howarth (1996) report evidence for a direct connection between wind features and an NR P mode. They found both in H$\alpha$ up to $0.3v_{\infty}$ and in He II $\lambda$5411 the same period of 8.54h. However, they consider the 19.6h period found in IUE observations incompatible with this NR P period.

The presence of multimodes, however, has possibly interesting consequences for the origin of cyclical wind variability (see also Rivinius et al. (1998) for the Be star $\mu$ Cen). Consider for example a case with two different prograde sectoral modes, traveling around the star with different frequencies. A given crest of the faster wave will at a certain moment overtake a crest of the slower wave, which means an enhancement of the total amplitude at a certain longitude. The next enhancement will be when a different crest will overtake, but this will be at a different longitude. This will give rise to cyclical surface amplitude enhancements which may cause wind perturbations that are related to the relative traveling speeds and the $m$ values of the NR P waves. The simultaneous presence of more than two modes will increase the complexity of this beating effect. Observations show, however, that the periods of cyclical wind variability of O (and B) stars scale with the rotation period of the stars (Pinjna 1988, Henrichs et al. 1988), and it is difficult to understand how this could be related to the above described effect of beating NR P modes.

We therefore think that the best candidate for the cause of the cyclical wind variability still remains the presence of weak magnetic fields on the surface, corotating with the star. A proof has to wait for a systematic deep survey of these fields. A preliminary upper limit of 70 G on the longitudinal component of the magnetic field strength of $\xi$ Per was presented by Henrichs et al. (1998a).

In conclusion, if our interpretation is correct, the number of confirmed O stars with NR P is now about 6 (see Fullerton et al. 1996, including $\zeta$ Pup). In the light of asteroseimological applications, we note that 5 of these are runaways. Although the statistics are limited, OB runaway stars tend to rotate rapidly and to have an enhanced surface He abundance (Blaauw 1993). These factors might play a role in exciting the NR P. The exception is the pulsator 10 Lac, which is not classified as a runaway star, but this star is associated with a bow shock detected on IRAS 60 micron maps (van Buren et al. 1995), indicating a post binary mass-transfer history, and hence a different internal structure is likely, perhaps favouring the existence of NR P modes.

Because of the known ubiquity of line-profile variables among O stars it one can expect that with a concentrated observational effort with sufficient S/N, coverage and time resolution, more pulsation-mode identifications in O stars are likely to follow. Higher spectral resolution is needed to find possible multimodes in these stars. The possible consequence of such multimodes for cyclical wind behavior needs to be investigated.

Acknowledgements. The help of skillful staffs at the Calar Alto and Kitt Peak observatories is highly appreciated. HFH thanks Rolf Kudritzki and his colleagues for their warm hospitality and encouragement during the period that this work was initiated, and for his suggestion to concentrate on the He I $\lambda$4713 line for searching for pulsations. We thank Mariano Méndez for drawing our attention to the minimum-entropy analysis and help. We are also grateful to the referee Alex Fullerton for his constructive comments. JAdJ acknowledges support by the Netherlands Foundation for Research in Astronomy (NFRA) with financial aid from the Netherlands Organization for Scientific Research (NWO) under project 781-71-053.

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Chapter 5

A search for the magnetic field of the O7.5III star ξ Persei, including an analysis of profile variability in multiple lines

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Abstract. We present circular spectropolarimetric observations of the O7.5 III star ξ Persei obtained during one week with the MusiCoS echelle spectropolarimeter mounted at the 2m Télescope Bernard Lyot to attempt to detect a surface magnetic field. Such a field is suspected to be the cause of the prominent cyclonical optical and UV spectral variability on the rotational timescale of this star, similar to other early-type stars. Only null-detections are obtained, with a 47 G 1σ error bar for the disc-averaged line-of-sight component of the surface magnetic field vector, from which we conclude that the polar strength of a dipole magnetic field structure is less than 500 to 800 G, depending on the assumed geometry.

These observations are also used to quantitatively analyse the profile variability in 8 optical lines by means of the use of Temporal Variance Spectra (TVS). The previously known 2.09 d wind period is recovered, whereas rapid redward moving features in a number of optical lines, in particular He II λ5411, are consistent with an earlier discovered non-radial ℓ=3 pulsation mode with a period of 3.5 h. We also report for the first time in this star blue-ward moving features in Hα recurring on about the same timescale as the NRP, which may indicate a direct impact of the pulsations on the wind.

5.1. Introduction

Like most O and many B stars, the O7.5 III star ξ Persei shows very prominent cyclonical wind variability in the UV resonance lines. This variability manifests itself in discrete absorption components (DACs) which migrate from red to blue, narrowing when they approach the terminal velocity (see e.g. Kaper et al. 1996, 1999a). In ξ Per the DAC period is 2.09 d (Henrichs et al. 1998a, de Jong et al. 2000a, Chapter 2). Multiwavelength observations of a number of OB stars have shown that the cyclic behavior is present down to the deepest layers of the wind, close to the surface of the star, and that the typical timescale varies with the rotational timescale. This strongly argues in favor of a surface phenomenon which perturbs the base of the flow.

In the so-called Corotating Interaction Region (CIR) model a perturbation at the surface of the star causes a local increase (or decrease) of the radiative force driving the stellar wind, resulting in a slow (or fast) stream in the stellar wind. Due to rotation of the star, the slow stream is overtaken by the faster undisturbed wind which leads to the formation of a spiral-shaped shocked region (the CIR), which corotates with the star. This model was proposed by Mullan (1984, 1986), whereas Cranmer & Owocki (1996) performed hydrodynamical computations, which successfully reproduce the development of the DACs in the UV.

Extensive observational campaigns carried out during the past decade have lead to support to the CIR model. For example, the wind modulations in the B0.5 Ib star HD 64760 are explained by a kinematic model based on CIRs (Owocki et al. 1995, Fullerton et al. 1997). Coordinated UV and optical (Hα) observations of ξ Persei (Kaper et al. 1997, Henrichs et al. 1998a, de Jong et al. 2000a, Chapter 2) are consistent with the CIR model. De Jong et al. (2000c, Chapter 7) modeled the wind variations in ξ Per based on the CIR picture. With a simple kinematic model, in which a spiral shaped density enhancement is rotated in a velocity field based on a β-law and conservation of angular momentum, they show that the CIR model very likely applies to the case of ξ Per.

In spite of all these efforts the origin of the perturbations leading to the formation of CIRs is still unknown. Unraveling the cause of the wind variability is one of the most challenging problems in this field of research of the past 15 years. The most likely candidates are magnetic fields and/or non-radial pulsations (NRPs). Weak magnetic fields perturb the gas pressure (or even confine the flow) and could form anchored surface structures that corotate with the star. NRPs will cause velocity and density perturbations at the base of the wind. The velocity with which a NRP bump moves with respect to the stellar surface depends on the NRP mode and frequency, and therefore the
timescale on which the wind is disturbed does not have to correspond with the rotation period. This could provide an observational test, but multimode beating could result in several timescales which cannot be calculated if not all NRP modes are known. In the review by Henrichs (1999b) only 6 confirmed pulsators are listed. This list includes ξ Per (de Jong et al. 1999a, Chapter 4) with an f=3 mode and a period of 3.45 h. The pattern velocity of this mode is 8.8 times the rotational velocity, which is difficult to comply with the wind period of 2.1 days. The same NRP mode has also been detected in observations of 1994 (de Jong et al. 2000a, Chapter 2) and 1996 (Henrichs et al. 1998b, Chapter 3). No other NRP modes are found.

The remaining alternative would be the presence of a surface magnetic field in O stars. None has been found so far in O-type stars. The star with the earliest spectral type and a magnetic field is β Cep B1 III, as was recently reported by Henrichs et al. (2000, Chapter 6). In this star the stellar wind is magnetically controlled, as can be seen from the different character of the UV profile variations as compared with the O-star DAC variability. We made several attempts to detect a feature in ξ Per (October 1994, Henrichs et al. 1998a, de Jong et al. 2000a, Chapter 2 and November 1996, Henrichs et al. 1998b, Chapter 3). In 1994 the Western Ontario Pockels cell polarimeter was used on the CFHT which uses the red and blue wings of the H/3 line (Landstreet 1982). During this observing run null-detections were obtained with 1σ error bars of 70 G. During the November 1996 MusiCoS campaign new magnetic field measurements were performed with the MusiCoS echelle spectropolarimeter (Donati et al. 1999a), also on the CFHT. Due to bad weather no improvements on the previous measurements could be achieved (Henrichs et al. 1998b, Chapter 3).

In this paper, we present new magnetic field measurements of ξ Per with the same MusiCoS polarimeter mounted on the 2.2 Télescope Bernard Lyot at Pic du Midi. During the same run the magnetic field of β Cep was discovered. In Sect. 5.2 we describe the instrument as well as the observing procedure. In Sect. 5.3 the observation log, data reduction procedure and magnetic field measurements are presented. No significant detection has been obtained. We derive an upper limit for the magnetic field in Sect. 5.4. We have also used the spectra for a detailed investigation of the profile variability in 8 optical lines. The analysis is presented in Sect. 5.5. We finally discuss our results in relation to the wind variability in Sect. 5.6.

5.2. Observations

We have obtained 38 circularly polarized (Stokes V) and unpolarized (Stokes I) spectra of ξ Persei, using the MusiCoS spectropolarimeter mounted on the 2 m Télescope Bernard Lyot (TBL) at the Observatoire du Pic du Midi (observers HPH and JAdJ). Our observing strategy aimed at (1) obtaining the most sensitive possible magnetic field measurements for this star and (2) observing the short-timescale lpv's simultaneously in many optical lines. The journal of observations is given in Table 5.1.

The spectropolarimetric setup consists of the MusiCoS fiber-fed cross-dispersed échelle spectrograph (Baudrand & Böhm 1992, Catala et al. 1993) with a dedicated polarimetric unit (described by Donati et al. 1999a) mounted at the Cassegrain focus. The light passes through a rotatable quarter wave plate after which the beam is split into two, along and perpendicular to the instrumental reference azimuth, respectively. Two fibers transport the light to the spectrograph, where both orthogonal polarisation states are simultaneously recorded. The spectral coverage in one exposure is from 450 to 660 nm with a resolving power of about 35,000. The Site CCD detector with 1024×1024 of 24μm pixels was used, which has a quantum efficiency exceeding 50% in the U band.

A complete Stokes V exposure consists of four subsequent subexposures between which the quarter wave plate is rotated by 90° back (called the q1 exposure) and forth (the q3 exposure), such that the two beams are exchanged throughout the whole instrument. The sequence of a complete exposure is q1→q3→q3→q1. With this procedure all systematic spurious circular polarisation signals down to 0.002% rms can be suppressed (Donati et al. 1997, Donati et al. 1999a). At the beginning and at the end of each night we took 15 flatfield exposures, whereas wavelength calibrations with a Th-Ar lamp were taken every hour, in addition to the usual bias frames and polarisation check exposures. For the reduction we used the flatfield series from the beginning of the night.

5.3. Data reduction and results

The data reduction was performed with the dedicated ES-PRIT reduction package, described by Donati et al (1997). With this package the geometry of the orders on the CCD is determined first, and after an automatic wavelength calibration based on the Th-Ar frames, a rigorous optimal extraction of the orders is performed. Besides combining the 4 subexposures in order to obtain the Stokes V profiles, we also extracted the subexposures separately. In this way we can study the spectrum at a time resolution of 15 minutes.

For the magnetic field measurements the profiles of 32 relatively weak lines, selected with a table appropriate for an O7.5 star, were combined by means of the Least Square Deconvolution (LSD) method (Donati et al. 1997) in the interval [-504, 562.5] km s⁻¹. This range was chosen such that an equal amount of continuum is available on both sides of the line profile, taking into account the radial velocity of 60 km s⁻¹. Many of these lines are blends, so that we are effectively left with 10 distinct lines. The noise in the LSD spectra is listed in Table 5.1, along with the signal to noise ratio obtained in the raw data. The LSD profiles were normalized outside the regions [-300, 300] km s⁻¹ and corrected for the radial and barycentric velocity of the star. An
A search for the magnetic field of the O7.5III star ξ Persei, including an analysis of profile variability in multiple lines

Table 5.1. Journal of observations of MuSiCoS polarimetry and the resulting magnetic field measurements of ξ Per at TBL at the Pic du Midi. The Barycentric Julian date is given at mid-exposure time (t_{exp}). Also listed is the quality of the Stokes V spectra, expressed as the S/N per 4.5 km s^{-1} around 550 nm in the raw spectrum, and the relative rms noise level N_{rSD} (per 4.5 km s^{-1} velocity bin) in the Least-Squares Deconvolved Stokes V spectra. The values for B_{l} and σ_{B_{l}} are listed in the last two columns.

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<td>120</td>
<td>0.403</td>
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Fig. 5.1. Example of a representative LSD stokes V (upper panel) and stokes I (lower panel) profile of ξ Per. The N (null) profile is shown mapped on LSD. The width of most profiles. Furthermore, the null profiles were examined, which are obtained by processing pairs of identical azimuths. Field measurements on these profiles reveal how much can be contributed by spurious signals due to instrumental effects. All the magnetic field measurements are listed in Table 5.1 and are plotted in Fig. 5.2.

5.4. An upper limit for the magnetic field

Among the measurements listed in Table 5.1 there is no significant field detection. With 71% of the B_{l} values deviating less than 1σ and 97% deviating less than 2σ the dataset is consistent with drawings from normal distributions with averages of 0 and standard deviations σ_{B_{l}}. The weighted averages of the three best nights (15/16, 16/17 and 17/18 December) are respectively \( \pm 40 \) G, \( \pm 47 \) G and \( \pm 38 \) G (1σ errors). From these values we conclude that the upper limit of our null detection is about 40 G. We cannot average several nights to improve on this upper limit, because the stellar rotation period is likely 4.18 d, which we derived from UV observations (see de Jong et al. 2000a, Chapter 2), and averaging over a shorter timescale would rather cause cancellation. For the sake of discussion we assume that the field varies with the EW as in β Cep. If the field varies with the well determined wind period of 2.1 d, our measurements could fall in the worst-case scenario at the zero points of the sinusoid. (The phase is not well enough determined to be predicted within 1 day.) New observations at intermediate phases could prove or disprove this. If the field varies with the rotation period of 4.2 d and if we consider that B_{l} was compatible with zero in our two best nights (MBJD=13.5 and 15.5) we can consider the error bar of the third best night in between as an upper limit, which means that \( B_{\text{max}} \approx 47 \) G. We note that if we again assume for the sake of discusion that the field varies with the EW as in β Cep, we would expect the maximum value of \( |B_{l}| \) at minimum EW, which falls around MBJD=14.5.
Although no significance can be attributed, the lowest value in the EW of Hα indeed coincides with the lowest value of \( B_x \), but more significant measurements need to be done to confirm this.

Our adopted upper limit of 47 G for the disc-average of the longitudinal component of the field allows for a much stronger polar field strength. To calculate the polar field strength one has to assume a model for the field configuration. An oblique dipole field seems the most likely model considering our argumentation that during every rotation of 4.18 d two DACs appear which slightly differ from each other (Kaper et al. 1996, de Jong et al. 2000a, Chapter 2).

With a tilt angle \( \beta \) between the magnetic dipole and the rotation axis, and given the inclination \( i \) and the rotational phase \( \phi \), the polar field \( B_p \) and \( B_x \) are related as follows (Preston 1967):

\[
B_x = B_p \frac{15+u}{20(3-u)} \times (\cos \beta \cos i + \sin \beta \sin i \cos 2\pi(\phi - \phi_0)),
\]

where \( u \) is the limb darkening parameter. According to Puls et al. (1996) the stellar parameters for \( \xi \) Per are

\[ T_{\text{eff}} = 36,000 \text{ K} \quad \text{and} \quad \log g = 3.4, \]

for which \( u = 0.22 \) can be derived (Wade & Rucinski 1985). The main uncertainties remain in \( \beta \) and \( i \). Adopting a rotation period of 4.2 d, and using a projected rotational velocity \( v \sin i = 204 \text{ km s}^{-1} \), an inclination of \( i \approx 40^\circ \) is implied for a radius \( R_* = 25R_\odot \) (Puls et al., 1996). In this case the poles of the magnetic field should be almost on the equator, because otherwise subsequent DACs would differ much more than observed (the pole on the facing hemisphere would be longer visible than the one on the other hemisphere). Therefore we assume \( \beta = 90^\circ \), which means that \( B_p = 5.7 B_{\text{max}} \) according to Eq. 5.2. For a significant detection \( B_{\text{max}} \) should be at least 3 times \( \sigma_B = 47 \text{ G} \), which gives an upper limit on \( B_p = 800 \text{ G} \) as a conservative estimate. If we assume \( \beta = 90^\circ \) Eq. 5.2 gives \( B_p \cos \beta = 3.65 B_x \), which gives for \( \beta = 90^\circ \) an upper limit of \( B_p = 515 \text{ G} \), but this is very uncertain considering the unknown geometry. So, depending on the assumed geometry the upper limit of \( B_p \) lies between 500 and 800 G.

5.5 Analysis of the line-profile variability

We have extracted 8 lines from the echelle spectra obtained in the subexposures. The line profiles were rebinned to 15 km s\(^{-1}\) bins using a triangular weighting function and normalized to the continuum using 2\(^{\text{nd}}\) or 3\(^{\text{rd}}\) order polynomials. The errors in the original spectra were consistently propagated. We had to remove ripples with an amplitude of about 10\(^{-3}\) of the continuum flux. These ripples are constant during a single night, so we could isolate the pattern from the night-averaged profile. This was done by dividing the average profile by its smoothed version. All spectra were subsequently divided by the ripple pattern. Some spectra still show ripple patterns (e.g. around MBJD=14.3) which we could not remove.

In order to make the variations better visible we divided each profile by the average profile of the whole dataset. The used lines and continuum regions are given in Table 5.2. We
A search for the magnetic field of the O7.5III star ξ Persei, including an analysis of profile variability in multiple lines

![Image](image.png)

**Table 5.2.** The lines which were extracted from the optical spectra. The element, rest wavelength, minimum velocity and order for the polynomial continuum fit are given. The transition from continuum to line is set at $-v_{\text{min}}$ and $+v_{\text{min}}$.

<table>
<thead>
<tr>
<th>Line</th>
<th>Wavelength (nm)</th>
<th>Continuum $v_{\text{min}}$ (km s$^{-1}$)</th>
<th>Order</th>
</tr>
</thead>
<tbody>
<tr>
<td>Hα</td>
<td>656.2797</td>
<td>700</td>
<td>2</td>
</tr>
<tr>
<td>Hβ</td>
<td>486.1323</td>
<td>700</td>
<td>3</td>
</tr>
<tr>
<td>He II Br γ</td>
<td>541.1516</td>
<td>350</td>
<td>3</td>
</tr>
<tr>
<td>He II Pa α</td>
<td>468.5698</td>
<td>350</td>
<td>3</td>
</tr>
<tr>
<td>He I</td>
<td>587.5614</td>
<td>350</td>
<td>3</td>
</tr>
<tr>
<td>He I</td>
<td>492.1931</td>
<td>350</td>
<td>3</td>
</tr>
<tr>
<td>He I</td>
<td>471.3143</td>
<td>350</td>
<td>3</td>
</tr>
<tr>
<td>O III</td>
<td>559.2252</td>
<td>350</td>
<td>3</td>
</tr>
</tbody>
</table>

5.5.1. Description of the variations

All the wind lines show variability on short and long timescales. In Hα and Hβ strong relative emission is visible around MBJD=2.4 and MBJD=4.4, which corresponds well with the known 2.1 d period (de Jong et al. 2000a, Chapter 2). There are also several varying features visible on a much shorter timescale: we can see blue-ward moving features with a separation of about 0.15 d in Hα around MBJD=1.4, which are shown better in Fig. 5.4.

Around MBJD=5.4 the He II Br γ, O III λ5592 and He I lines all show red-ward moving features with a recurrence time of about 0.15 days. Figure 5.7 shows that these variations are likely due to the same NRP mode as observed in October 1989 (de Jong et al. 1999a, Chapter 4). In this figure the phase of this NRP with $P=3.45$ h and $f=3$ is overplotted over our He II Br γ dynamic spectrum. Possibly, another mode may be present considering the red-ward moving features around MBJD=3.5, with a recurrence time of about 0.1 days.

5.5.2. Temporal Variance Spectra

We applied the Temporal Variance Spectrum (TVS) method (Fullerton et al. 1996) to our data in order to quantify the amount of variability in each line. It was possible to apply this method fully on our data, since the statistical noise values are provided by EsPRIT. The normalized dynamic spectrum is defined in a matrix $S$ where element $S_{ij}$ represents the $j$th wavelength bin of the $i$th spectrum. We define $\sigma_{ij}$ as the error of this bin and $\sigma_{ic}$ as the average error in a continuum bin of spectrum $i$. The TVS is defined as follows:

$$\text{TVS}_j = \frac{1}{N-1} \sum_{i=1}^{N} \left( \frac{\sigma_{ij}}{\sigma_{ic}} \right)^2 \left[ S_{ij} - \bar{S}_j \right]^2$$

(5.3)

Here $\sigma_0$ is a standard continuum noise level which has the following value:

$$\sigma_0 = \left( \frac{1}{N} \sum_{i=1}^{N} \sigma_{ic}^{-2} \right)^{-1}$$

(5.4)

and $\bar{S}_j$ is the weighted average line profile:

$$\bar{S}_j = \frac{1}{N} \sum_{i=1}^{N} \left( \frac{\sigma_0}{\sigma_{ic}} \right) S_{ij}$$

(5.5)

From now on we will plot and use only $\sqrt{\text{TVS}}$, because this value scales linearly with the size of the spectral deviations. It can best be interpreted as the average amplitude of the variations in each bin. If there are no other variations
than noise we have $\sqrt{TVS} \sim \sigma_0$. Figure 5.8 shows the TVS for all 8 lines. In order to quantify the total amount of variability in each line we defined a “TVS Equivalent width” analogous to the normal Equivalent Width of a profile:

$$EW_{TVS} = \sum_{j=1}^{N} \sqrt{TVS_j - TVS_{cont}}$$

which can be compared with the EW of the average line profile:

$$EW = \sum_{j=1}^{N} 1 - \bar{S}_j$$

Both values were computed for all lines between $-400$ and $+400$ km/s and are given in Table 5.3. The errors in the EW were computed following Chalabaev & Maillard (1983). Most lines seem to have about the same amount of variability relative to their EW. We find $EW_{TVS}/EW \sim 4 - 5\%$. H$\alpha$ and He I Pa $\alpha$ have a high relative variability of 13% and 10%. Both these lines have a relatively strong contribution from the stellar wind and belong to comparable atomic transitions.

5.5.3. Period analysis

In order to quantify the timescale(s) of variability, we performed a period analysis using the CLEANed Discrete Fourier Transformation (CDFT, Roberts et al. 1987). We searched for the known 2.09 d (de Jong et al. 2000a, Chapter 2) period and NRP period of 3.45 d (de Jong et al. 1999a, Chapter 4). The periodogram of H$\alpha$ (Fig. 5.9) shows a main period at 2.06(4) d between [−350, −200] and [0, 450] km s$^{-1}$. This means that the wind variability has the same periodicity as in previous years (1991, 1994 and 1996). However, the variability in the blue wing is much less than in 1994 and the power at this period is dominated by the red wing, like in 1987 (Kaper et al. 1997). In spite of the clear presence of NRP patterns in the dynamic spectra none of the CDFT’s of the photospheric lines and He I Br $\gamma$ revealed any significant peak above 1 c/d.

5.6. Discussion and conclusions

The upper limit of about 500 G on the magnetic field of $\xi$ Per leaves enough room for a field which can perturb the wind strongly enough to cause the observed wind variability. By assuming that the magnetic field causes a significant dark spot (when the magnetic pressure equals the gas pressure) we estimate that a field of 100 G should be suf-
A search for the magnetic field of the O7.5III star $\xi$ Persei, including an analysis of profile variability in multiple lines

Table 5.3. Total amount of variability compared to the EW. These values were computed using Eq. 5.6 and 5.7.

<table>
<thead>
<tr>
<th>Line</th>
<th>EW (km s$^{-1}$)</th>
<th>EW$_{TVS}$ (km s$^{-1}$)</th>
<th>EW$_{TVS}$/EW</th>
</tr>
</thead>
<tbody>
<tr>
<td>H$\alpha$</td>
<td>77.28(8)</td>
<td>10.50(6)</td>
<td>0.1358(8)</td>
</tr>
<tr>
<td>H$\beta$</td>
<td>119.1(1)</td>
<td>5.23(3)</td>
<td>0.0440(3)</td>
</tr>
<tr>
<td>He II Br $\gamma$</td>
<td>43.0(1)</td>
<td>1.85(2)</td>
<td>0.0430(6)</td>
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<td>He II Pa $\alpha$</td>
<td>23.3(2)</td>
<td>2.45(3)</td>
<td>0.105(2)</td>
</tr>
<tr>
<td>He I $\lambda$5875</td>
<td>57.24(7)</td>
<td>3.33(4)</td>
<td>0.0582(7)</td>
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<tr>
<td>He I $\lambda$4922</td>
<td>20.75(7)</td>
<td>0.78(2)</td>
<td>0.038(1)</td>
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<tr>
<td>He I $\lambda$4713</td>
<td>14.26(8)</td>
<td>0.54(3)</td>
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<tr>
<td>O III</td>
<td>16.40(7)</td>
<td>0.86(2)</td>
<td>0.052(1)</td>
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Fig. 5.9. Greyscale representation of the periodogram as a function of velocity of the H$\alpha$ spectra. The side panel shows the power summed over all velocities. The highest peak is normalized to unity. The top panel displays the ratio of the observed standard deviation to the expected standard deviation.

Acknowledgements. The helpful assistance of the observatory staff members at TBL is well remembered and greatly acknowledged. We thank Rens Waters for his helpful comments. JDJ acknowledges support by the Netherlands Foundation for Research in Astronomy (NFRA) with financial aid from the Netherlands Organization for Scientific Research (NWO) under project 781-71-053.

References

...some Ap/Bp stars. The cyclical variability in this star with a period of 15.4 days and the correlation with the X-ray variability are similar to that of magnetic chemically peculiar stars (see e.g. Wade, 1997). Sofar, several attempts to measure the field in this star have been unsuccessful (Donati & Wade 1999b, Mathys 1999). This star is 2.5 times fainter than $\xi$ Per, but its relatively low $v\sin i \simeq 25$ km s$^{-1}$ largely compensates for this in the S/N of the $B_t$ measurements. It is not clear whether the MCWS can be applied to $\xi$ Per, taking into account the different spectral type and environment. Depending on which model applies to $\xi$ Per we need to decrease the error bars on our magnetic field measurements by about 2 to 5 times to be able to measure such a weak field. This means that we must collect 4 to 25 times more light, which should be feasible with 4 to 8 m-class telescopes.

The line-profile variations in the 8 dynamical spectra of the optical lines show some interesting variations. Besides the 2.09 d period which indicates that $\xi$ Per still shows the usual wind variability also much variations are seen on much shorter timescales. In particular H$\alpha$ shows several narrow bluward moving features around MBJD=1.4 which seem to recur on the same timescale as the NRP features in the photospheric lines around MBJD=3.5 (see Fig. 5.4). Furthermore, the emission in H$\beta$ around MBJD=5.4 has some bright spots which are in phase with the NRP pattern in He II Br $\gamma$. These variations in H$\alpha$ and H$\beta$ may well represent the dynamic response of the stellar wind to the NRP. They cannot just be the NRP themselves since they move blue-ward and beyond the projected rotational velocity of 200 km s$^{-1}$. New investigations of this phenomenon is clearly needed. We are continuing our search for the photospheric origin of wind variability.

References
Fig. 5.5. Greyscale images of the dynamic spectra of Hα, Hβ, He I λ5876 and He II Pa α. The uppermost panels show the average line profiles. The lower panels show the quotient spectra.
Fig. 5.6. Greyscale images of the dynamic spectra of He II Brγ, O III λ5592, He I λ4922 and He I λ4713, in the same format as Fig. 5.5.
Chapter 6

The magnetic field of $\beta$ Cephei – a key system towards understanding the Be phenomenon?

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To be submitted to A&A

Abstract. New time-resolved circular spectropolarimetric observations of the B1 IIIe star $\beta$ Cep ($v\sin i = 27$ km s$^{-1}$) obtained with the MuSiCoS echelle spectropolarimeter and the 2 m Télescope Bernard Lyot show a sinusoidally varying magnetic field with a strength between +90 G and −90 G for the averaged line-of-sight component. The period corresponds very accurately with the 12 day period as derived from stellar wind variations observed in the ultraviolet. The epoch of the positive maximum field corresponds in phase with the maximum emission in the UV wind lines. This gives compelling evidence for a magnetic-rotator model for this star, with an unambiguous rotation period of 12 days.

The relative phasing of the magnetic and UV line measurements are discussed, and preliminary modeling of the circumstellar envelope of this star as a magnetically-confined wind shock predict an X-ray flux which is in agreement with the observed value.

We discuss the magnetic behavior as a function of pulsation behavior, UV line variability and H$\alpha$ episodes. The observed radial velocities are in agreement with the predicted values near periastron passage around its binary companion with 90 y period. We propose that the presently slow rotation of the star is due to magnetic braking.

The similarity between the intermittent H$\alpha$ emission phases in $\beta$ Cep and in Be stars suggests that the primary origin of the Be phenomenon does not have to be rapid rotation: in $\beta$ Cep the necessarily high velocity to bring material in (keplerian) orbit is provided by the high corotation velocity at the Alfvén radius (about 10 stellar radii), whereas in the case of rapidly rotating Be stars the rotation of the surface itself could be sufficient. In both cases the cause of the emission phases has still to be found, but it is noted that the occurrence of temporary magnetic fields, much weaker than detected here, remains the strongest candidate.

This is the first confirmed detection of magnetic field in a non-chemically peculiar upper main-sequence star.

Key words: B stars – Be stars – magnetic fields – pulsation – winds – binaries

6.1. Introduction

Besides the well-studied pulsation period of 4$^h$34$^m$ (Telling et al. 1997), the star $\beta$ Cep (HD 205021, spectral type B1 III, Lesh 1968, but earlier references give also B2 III or B1 IV) shows a very significant period of 12 d in the equivalent width of the ultraviolet resonance lines. At its discovery by Fishel and Sparks (1972) with the OAO-2 satellite there was still an ambiguity between 6 and 12 days, but later investigations with IUE data (Henrichs et al. 1993, Henrichs et al. 1998a) left no doubt that the two minima in equivalent width of the C IV stellar wind lines, which are separated by 6 d, are unequal, and that the real period is 12 d. Henrichs et al. (1993) proposed that the UV periodicity arises from the 12 d rotational period of the star and suggested that the stellar wind is modulated by an oblique dipolar magnetic field at the surface.
Support for this hypothesis was given by the striking similarity between the UV-line behavior of $\beta$ Cep and of known rotating magnetic B stars, for example the B2 V helium-strong star HD 184927 (Barker et al. 1982, Wade et al. 1997), and by the reported (but not confirmed) average magnetic field strength of $B = (810 \pm 170) \, \text{G}$ for $\beta$ Cep itself by Rudy and Kemp (1978). A rotational period of 12 days corresponds well with an adopted radius between 6 and 10 solar radii, given the reported values of 20 – 43 km s$^{-1}$ for vsini.

To verify this hypothesis Henrichs et al. (1993) presented new magnetic field measurements obtained with the University of Western Ontario photoelectric Pockels cell polarimeter and 1.2m telescope, simultaneously with UV spectroscopy with the IUE satellite. They measured the magnetic field by polarimetry in the H$\beta$ line (Landstreet 1982 and references therein). The 12 day UV period in the equivalent width of the stellar wind lines of CIV, SiIII, SiIV and NV was confirmed, but much lower values for the magnetic field were found, with 1σ error bars of about 150 G, comparable to the measured field strength. Additional magnetic measurements with the same instrumentation by G. Hill (private communication) could not confirm the 12 d period, which was unexplained. It also remained puzzling why these new magnetic field measurements showed a much lower field than in 1987. The suggestion was put forward that perhaps the new Be phase of the star, discovered in July 1990 by Mathias et al. (1991), see also Kaper et al. (1992) and Kaper & Mathias (1995), might have been related to the decrease in magnetic field strength, but this could not be tested.

These considerations motivated us to undertake new magnetic measurements of $\beta$ Cep with the much more sensitive MuSiCoS polarimeter at the Pic du Midi observatory in France. Using this instrument has the clear advantage that all available (mostly metallic) lines can be selected in the spectrum, rather than just one Balmer line, which may be contaminated with some emission. In section 6.2 we describe the experimental setup and the observations. Section 6.3 summarizes the reduction and the results. In section 6.4 we compare the known periodicities in $\beta$ Cep with the magnetic measurements. In section 6.5 we pay attention to the H$\alpha$ emission phase of the star. In section 6.6 the relative phasing of the magnetic and UV line measurements are discussed, and we describe preliminary modeling of the circumstellar envelope of this star as a magnetically-confined wind shock. In the last section we give our conclusions and discuss the implications of the current measurements.

### 6.2. Observations

During December 1998, January, June and July 1999 we have obtained 23 circularly polarized (Stokes $V$) and unpolarized spectra (Stokes $I$) of $\beta$ Cep, using the MuSiCoS spectropolarimeter mounted on the 2 m Télescope Bernard Lyot (TBL) at the Observatoire du Pic du Midi (observers HFH, JADJ, GAW, SLSS, AT and EV). The strategy of the observations was threefold: (1) to cover the known 12 d period of the UV lines, (2) to have reasonable coverage during one pulsational period of 4.6 h, and (3) to study the behavior at a half-year timescale. The journal of observations is given in Table 6.1.

The spectropolarimetric setup consists of the MuSiCoS fiber-fed cross dispersed echelle spectrograph (Baudrand & Böhm 1992, Catala et al. 1993) with a dedicated polarimetric unit (described by Donati et al. 1999a) mounted at the Cassegrain focus. The light passes through a rotatable quarter wave plate after which the beam is split into two beams, along and perpendicular to the instrumental reference azimuth, respectively. Two fibers transport the light to the spectrograph, where both orthogonal polarisation states are simultaneously recorded. The spectral coverage in one exposure is from 450 to 660 nm with a resolving power of about 35000. The Site CCD detector with 1024×1024 of 24μm pixels was used, which has a quantum efficiency exceeding 50% in the U band.

A complete Stokes $V$ exposure consists of four subsequent subexposures between which the quarter wave plate is rotated by 90° back (called the q1 exposure) and forth (the q3 exposure), such that the two beams are exchanged.

<table>
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<th>S/N</th>
<th>$N_{\text{LS}}$</th>
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<td>0.015</td>
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</table>
| 23  | 1999 July 7 | 217.4174       | 60               | 790 | 0.014        

Table 6.1: Journal of observations of MuSiCoS spectropolarimetry of $\beta$ Cep at TBL at the Pic du Midi. The Barycentric Julian Date is given at mid-exposure time ($t_{\text{exp}}$). Also listed is the quality of the Stokes $V$ spectra, expressed as the S/N per 4.5 km s$^{-1}$ around 550 nm in the raw spectrum, and the relative rms noise level $N_{\text{LS}}$ (per 4.5 km s$^{-1}$ velocity bin) in the Least-Squares Deconvolved Stokes $V$ spectra.
The magnetic field of \(\beta\) Cephei – a key system towards understanding the Be phenomenon?

Throughout the whole instrument, the sequence of a complete exposure is \(q1 - q3 - q3 - q1\). With this procedure all systematic spurious circular polarisation signals down to 0.002\% rms can be suppressed (Donati et al. 1997, Donati et al. 1999a). At the beginning and at the end of each night we took 15 flatfield exposures, whereas wavelength calibrations with a Th-Ar lamp were taken every hour, in addition to the usual bias frames and polarisation check exposures. For the reduction we used the flatfield series nearest in time to the observations.

6.3. Data reduction and results

The data reduction was done with the dedicated ESpRIT reduction package, described by Donati et al. (1997). With this package the geometry of the orders on the CCD is first determined, and after an automatic wavelength calibration on the Th-Ar frames, a rigorous optimal extraction of the orders is performed. Finally the profiles of 63 relatively weak lines, selected with a table appropriate for early B stars, were combined by means of the Least Square Deconvolution (LSD) method in the interval \([-243, 243]\) km s\(^{-1}\). Many of these lines are blends of multiplets, leaving effectively 30 distinct lines. The noise in the LSD spectra was measured and given in Table 6.1, along with the signal to noise ratio obtained in the raw data. The profiles were normalized outside the regions \([-120, 120]\) km s\(^{-1}\).

Because the radial velocity amplitude as a consequence of the pulsation of \(\beta\) Cep is considerable, we shifted the minima of the \(I\) profiles (at \(v_{\text{min}}\), determined by a parabola fit) to zero velocity before calculating the longitudinal field strength. The profiles are often asymmetric, implying that the minimum flux does not occur at the radial velocity \(v_{\text{rad}}\) of the star, which was measured using the first moment of the profile with respect to the barycentric restframe, normalized by the equivalent width (we followed Schrijvers et al. 1997). The measured values of \(v_{\text{min}}\) and \(v_{\text{rad}}\) are included in Table 6.2. For typical examples of Zeeman LSD profiles (Stokes \(I\) and \(V\) in velocity space), see Fig. 6.1 for a zero, positive and negative field, respectively, the latter at two different pulsation phases of the star (see also below).

To compute a value for the mean longitudinal field \((B_t)\) from the LSD spectra, we took the first-order moment using the relation (Mathys 1989, Donati et al. 1997):

\[
B_t = \frac{-2.14 \times 10^{11} G}{\lambda c} \frac{\int v V(v) dv}{\lambda c \int [1 - I(v)] dv},
\]

(6.1)

where \(\lambda\), in nm, is the mean wavelength, \(c\) is the velocity of light (in the same units as \(v\)), and \(g\) is the mean value of the Landé factors of all lines used to construct the LSD profile. We used \(\lambda = 512.5\) nm and \(g = 1.234\). We approximated the integral in Eq. 6.1 by a simple summation in a range between \(\pm v_{\text{limit}}\) and computed the uncertainties as follows:

\[
\sigma_B = |B_t| \sqrt{\frac{\sum \sigma_{I_i}^2}{(\sum 1 - I_i)^2} + \frac{\sum \sigma_{V_i}^2 \sigma_{V_i}^2}{(\sum v_i V_i)^2}},
\]

(6.2)

The limits of the integral in Eq. 6.1 were carefully determined in order to minimize the uncertainties. We first
Ccmath has been calculated with the ephemeris given by Pigulski & phases of the radial velocity curve (with phase 0 defined at maximum) has been calculated with the ephemeris given by Pigulski & Boratyn (1992) in column 3. The measured radial velocity (accuracy: 2.5 km s\(^{-1}\)) is given in column 4, whereas the velocity shift, measured at minimum flux, used before calculating the magnetic field is given in column 5. The last two columns give the magnetic field values and their 1-\(\sigma\) uncertainties.

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<th>(V_{\text{min}}) km s(^{-1})</th>
<th>(B_t) G</th>
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The measured values are plotted as a function of time. The variability is obvious. We overplotted the best-fit sine function with parameters derived below.

### 6.4. Period analysis

The two main periodicities in \(\beta\) Cep are \(4^h34^m\) of the radial pulsation mode, studied since 1902, and 12 d in the absorption in the UV wind lines, known since 1972. We investigate whether the observed magnetic variability is related to these periods.
6.4.1. Pulsation period

The measured radial velocities of the star are given in Table 6.2. For the calculation of the phase in heliocentric radial velocity due to the radial mode of the pulsation we used the ephemeris for the expected maximum from Pigulski & Boratyn (1992) with $P = 0.1904852$ d and $T_{\text{max}} = 2413499.5407$ (column 3 in Table 6.2)

In Fig. 6.3 we plotted the measured radial velocity together with the magnetic field strength as a function of the calculated pulsation phase. From the figure it is clear that there is no correlation between the pulsation phase and the magnetic field, as expected.

![Fig. 6.3. Measured radial velocity (filled symbols, scale on the left) and magnetic field strength (open symbols, scale on the right) as a function of pulsation phase. As expected, no correlation between the two quantities is present in the data: at several occasions very different magnetic values are measured at a given pulsation phase. The observed system velocity as well as the difference between the observed and calculated phase of the maximum radial velocity confirm the predicted values for the star in its binary orbit near periastron passage](image-url)

A fit of a simple sinusoid through the radial velocity data with the known period kept fixed yields 19(1) km s$^{-1}$ for the semi amplitude, $-20(1)$ km s$^{-1}$ for the average system velocity and a maximum velocity at $\phi = 0.36(1)$. The phase shift can be compared with the difference between the observed and calculated (O-C) radial velocity maximum values predicted by Pigulski & Boratyn, which is used by them to calculate the binary orbit. With this ephemeris, our observation # 10 is taken during pulsation cycle 197741. Our phase difference of $-0.64(1)$, corresponding to $-0.112(2)$ d, is in good agreement with the expected phase delay caused by the light-time effect in the binary orbit near periastron (Pigulski & Boratyn 1992, Fig. 1). The system velocity of $-20(1)$ km s$^{-1}$ is also in good agreement with the extrapolated value (see their Fig. 4). Our observed values are in fact very close to the predicted values at periastron itself. Adopting that our data were taken at periastron would imply from their figures that the binary period is 85 y (from the phase delay) or 88 y (from the system velocity), in good agreement with the derived orbit of 91.6±3.7 y derived in 1990. New speckle interferometry may confirm the orbital phase.

6.4.2. UV stellar wind period

From IUE spectra it is known that the UV stellar wind lines of C IV, Si III, Si IV, N V and Al III show a very clear 12 day periodicity. In Fig. 6.4 we show representative C IV doublet UV profiles (from Henrichs et al. 1998a). Note that the outflow velocity exceeds $-600$ km s$^{-1}$. We also note that this type of variability is very unlike what is observed in O stars (e.g. Kaper et al. 1996), but is very similar to profile variations of other magnetic B stars (Henrichs et al. 1998a).

We measured the equivalent-width (EW) of this line in the velocity range $[-680, 945]$ km s$^{-1}$ after normalizing all 88 available IUE spectra between 1979 and 1995 to the same continuum around the C IV line and dividing each spectrum by the average of the normalized spectra. The error bars are calculated following Chalabaev & Maillard (1983), where we used a noise model for the IUE spectra according to Henrichs et al. (1994) with parameters $A = 27.3$ and $B = 873$. The resulting EW values are plotted as a function of phase in Fig. 6.5 (upper panel), in which we separated the early and most recent UV data to visualize the compatibility. We derived the following ephemeris for the deepest minimum, i.e. maximum emission:

$$T(\text{EW}_{\text{min}}) = \text{BJD 2445621.722(1)} + N \times 12.00106(6)$$

with $N$ the number of cycles. This period was derived from a Fourier analysis and sine fits, where the three highest peaks in the Fourier power diagram were used as starting conditions for the sine fits. The solution with the lowest reduced $\chi^2(4)$ was found at 6.0005313 d. We finally multiplied this period by 2 because of the significantly unequal two minima. The epoch of the deepest minimum is determined by a parabola fit through the relevant points after rephasing with the parameters of the sine fit. The uncertainty is derived with a Monte Carlo method with 1000 iterations (see de Jong et al. 1999a, Chapter 4 for a description). It looks if the maxima are not exactly separated by 0.25 in phase from the minimum, which will put constraints on modelling the outflow.

6.4.3. Magnetic period

We fitted a cosine function of the following form to the 23 $B_l$ measurements:

$$B_l(t) = B_0 + B_{\text{max}} \cos \left( 2\pi \left( \frac{t - t_{\text{av}}}{12.00106} + \phi \right) \right)$$

(6.4)

in which $t_{\text{av}}$ is the average of the times of observations. The fitted values are $B_0 = 0 \pm 5$ G, $B_{\text{max}} = 90 \pm 6$ G, $t_{\text{av}} = 2451239.457$ and $\phi = -0.093(10)$ with a reduced $\chi^2 = 0.75$. The quoted 1-$\sigma$ errors on the parameters are identical using a standard Levenberg-Marquardt method (Press et al. 1992).
Chapter 6

Fig. 6.4. Representative CIV profiles from IUE spectra showing the typical variation over a 12 day cycle, very similar to the type of variation observed in other magnetic B stars, but unlike the variations observed in O stars. The two doublet rest wavelengths are indicated.

Fig. 6.5. Upper panel: Equivalent width of the CIV stellar wind line measured in IUE spectra taken during 16 years as a function of phase calculated with Eq. 6.3. The deepest minimum at phase 0 corresponds to the maximum emission. Lower panel: Magnetic data as a function of the UV phase. The dashed sine curve has the same parameters as in Fig. 6.2. Note that there is no significant difference in zero phase between the UV and magnetic data and that the field crosses zero at the EW maxima.

or using the same Monte Carlo method as above. With the derived phase we find for the ephemeris of the maximum value of the field strength:

\[ T(B_{\text{max}}) = \text{BJD} 2451238.34(12) + N \times 12.00106 \ (6.5) \]

In Fig. 6.2 we have drawn a sine wave with this period and phase through the magnetic measurements. A comparison with the phase of the UV data (Eq. 6.3) shows that a deep EW minimum is predicted at BJD 2451238.22(3), which is, within the uncertainties, identical to the phase of maximum (positive) magnetic field. In Fig. 6.5 (lower panel) we have drawn a sine wave with the same parameters as used in Fig. 6.2 through the values of the magnetic field strength, and phased with the UV period. It is clear from the figure that the phase of minima of the stellar wind absorption (i.e. maximum emission) coincides very well with the extremes of the magnetic field, and that the maximum wind absorption coincides with field strength zero.

It is interesting to note that \( B_0 = 0 \pm 5 \) G implies that the asymmetry with respect to zero must be very small. The fact that the second EW minimum is shallower will put constraints on the geometry of the field. It is clear, however, that more magnetic data are needed to confirm the absence of asymmetry.

6.5. Hα behavior

Since \( \beta \) Cep is by the current definition a Be star (see for earlier Hα emission-phase histories Mathias et al. 1991 and Panko & Tarasov 1997), it is important to investigate the Hα emission state and its possible relation to the reported magnetic field strength. During our observations of \( \beta \) Cep there was no emission in Hα. In Fig. 6.6 we compare our results with earlier data taken at the Observatoire de Haute Provence (OHP). The emission phase has apparently declined beyond detectability since 1990, in accordance with the linear decay rate predicted by Kaper & Mathias (1995). The question arises whether during future emission phases the magnetic field structure will remain unaltered.

6.6. Plasma environment of \( \beta \) Cep

6.6.1. A magnetically confined disk

We describe here how the relative phasing of the magnetic and UV line measurements suggest that they are qualitatively consistent with the presence of a warm, moderately dense disc confined to the magnetic equatorial plane (e.g. Shore & Brown 1990, Babel & Montmerle 1997a).

The variation of the magnetic field and UV lines according to the same period (presumably the rotational period of the star) provides strong confirmation that the UV line variations of \( \beta \) Cep are due to the same phenomenon responsible for similar variability in the He-peculiar stars: the presence of a structured wind and magnetosphere which is trapped in the quasi-dipolar magnetic field and forced to
The magnetic field of \( \beta \) Cephei – a key system towards understanding the Be phenomenon

6.6.2. A magnetically confined wind shock?

We finish this section by describing preliminary modeling of the circumstellar envelope of this star as a magnetically-confined wind shock (Babel & Montmerle 1997a).

Babel & Montmerle (1997a) described a new theoretical model (the Magnetically-Confined Wind Shock, or MCWS, model) describing the formation of a structured magnetosphere around a magnetic early-type star with radiatively-driven wind. The MCWS model, based largely on earlier work by Havnes & Goertz (1984), was originally developed to explain the X-ray emission of some Ap/Bp stars (Babel & Montmerle 1997a), although it has been successfully applied to the young O7 star \( \theta^1 \) Ori C as well (Babel & Montmerle 1997b). In this model, a dipolar magnetic field confines and directs the stellar wind in the two magnetic hemispheres toward the (magnetic) equatorial plane, where their collision produces a strong shock. The resulting circumstellar structure consists of a higher-density, corotating equatorial “cooling disc”, and a high-temperature, lower-density shock and extended postshock region (Babel & Montmerle 1997a). The X-ray emission computed from the model is largely dependent on the assumed magnetic field.

The qualitative similarities between the observed and inferred properties of \( \beta \) Cep and the predictions of the MCWS model motivated us to investigate their quantitative agreement. For reasonable values of the mass (\( M = 15 \, M_\odot \)), effective temperature (\( T_{\text{eff}} = 23000 \rightarrow 25000 \, K \)), mass loss rate (\( M = 10^{-10} \, M_\odot \, \text{yr}^{-1} \)) and wind speed (\( v_w = 500 \, \text{km s}^{-1} \)), an MCWS can reproduce both the X-ray luminosity and temperature (\( \log L_x = 29.81 \) and \( T_x = 0.24 \, \text{keV} \), Berghoefer, Schmitt & Cassinelli 1996). This suggests that the wind and magnetosphere of \( \beta \) Cep may be well-described by such scenario.

6.7. Conclusions and discussion

We have found unambiguously a varying “weak” magnetic field in \( \beta \) Cep, consistent with an oblique dipolar magnetic rotator model (rotation period 12 d), in which outflow (causing wind emission) occurs along the magnetic poles, similar to models by Brown et al. (1985), Shore (1987), Shore et al. (1987), Shore & Brown (1990) and Babel & Montmerle (1997a). In these models the \( C^+ \) production in the jet-like mass loss is presumably due to the dissipation of shear-generated Alfvén waves near the polar cones.

For a dipolar field, the ratio of the values at the magnetic extremes \( r = B_{\text{min}}/B_{\text{max}} \) is related to the inclination angle, \( i \), and the angle of the magnetic axis with respect to the rotation axis, \( \beta : r = \cos(\beta + i)/\cos(\beta - i) \) (Preston 1967). We found above that there is no significant asymmetry between the magnetic extremes, which is consistent with an equator-on inclination angle, and with the magnetic axis perpendicular to the rotation axis. If we adopt an inclination angle of 60° and a projected rotational velocity of 27(3) km s\(^{-1}\) (as derived by Teltting et al. 1997 from a pulsation mode analysis), the rotation period requires a radius of 7.4(8) \( R_\odot \), which is in rather good agreement with the well-determined radius of the B1 III star \( \epsilon \) Cen of 5.9\(^{+1.2}_{-0.6}\) \( R_\odot \) based on interferometric and parallax measurements.

We emphasize that we have only measured the longitudinal component of the magnetic field, i.e. the component in the line of sight of the averaged value over the stellar disk. The intensity at the magnetic poles must be stronger. For
a perpendicular magnetic rotator the polar field of a dipole is \(3.2 \times 10^{10}\) \( \text{G} \) (Schwarzschild 1950), i.e. about 300 \( \text{G} \). The fact, however, that the EW curve in the stellar wind lines has two unequal maxima at epochs when the projected field is strongest, suggests that there should be a slight asymmetry present. This could of course be due to a slightly different geometry (off-centered dipole, or higher-order fields) at the two hemispheres, which can easily be hidden in the observed field strength which is the integrated value over the visible surface.

We note that a configuration as found here will cause magnetic braking, and this might explain the low rotational velocity of this (evolved) Be star. The spindown timescale of a star rotating with angular velocity \(\Omega\) is equal to \(\tau_{\text{sd}} = \frac{\tau}{\Omega}\). The rate of change, \(\frac{\tau}{\Omega}\), is caused by the amount of angular momentum carried away by the corotating stellar wind with mass-loss rate \(\dot{M}\) at the Alfvén radius \(r_A\), and can be calculated from

\[
\dot{M}v_w r_A = I \dot{\Omega}.
\]

The Alfvén radius, scaled with the stellar radius, \(R_*\), as a function of the stellar wind parameters is given by

\[
r_A \approx 12.5 R_* \frac{B^{1/2}}{B_{\text{Sun}}} \left( \frac{v_\infty}{v_\text{Sun}} \right)^{1/2} \left( \frac{M_*/10^{-10}}{\Omega_{\text{Sun}}^{-1}} \right)^{1/4}
\]

where the polar field strength \(B\) is in units of 300 \(\text{G}\), the wind velocity \(v_{\infty}\) is given in units of 600 \(\text{km s}^{-1}\), and the mass-loss rate \(\dot{M}\) is in units of \(10^{-10} \text{M}_\odot \text{y}^{-1}\). For the typical parameters of \(\beta\) Cep given above the spindown timescale is \(1.6 \times 10^7\) \(\text{y}\), which is shorter than the main-sequence timescale of a 15 \(\text{M}_\odot\) star. This is a conservative estimate, because the field was assumed to be constant, whereas on the main sequence the star most likely had a stronger field, at least because of its smaller radius. This short spindown timescale gave the star sufficient time during its course to its present slightly evolved stage to slow its rotation rate down to its low value, nearly independent of its original rotation speed. This conclusion has an interesting consequence in the light of the well documented \(\text{H}\alpha\) emission history of \(\beta\) Cep, which is very similar to that of the usually rapidly rotating Be stars. The (unknown) primary mechanism causing the Be phenomenon, i.e. hydrogen-emission episodes on a timescale of years to decades, is apparently the same in the slow rotator \(\beta\) Cep. The role of the rapid rotation in Be stars is to give the temporary excess of outflowing matter sufficient angular momentum to stay in a keplerian disk around the star. In \(\beta\) Cep such matter would leave the star not with the angular momentum corresponding to the rotation velocity at the surface but to the corotating velocity at the Alfvén radius, which is more than 10 times larger, and very much comparable to the surface rotation velocity of rapidly rotating Be stars. This suggests that the primary origin of the Be phenomenon is not necessarily rapid rotation. We note that temporarily emerging magnetic field structures, smaller than measured here, could cause the outflowing matter to go into orbit, and may be responsible for the Be phenomenon, but this hypothesis cannot be disproven as yet because of instrumental limitations. Spindown will only occur when the magnetic pressure is stronger than the outflowing pressure of the wind. This implies for a star like \(\beta\) Cep a field of at least 15 \(\text{G}\), with a spindown timescale of \(5 \times 10^7\) \(\text{y}\). Such small fields cannot be presently detected, but should be searched for in future investigations.

The role of variable pulsation behavior with respect to the hydrogen emission episodes, if such a role exists, is probably different from rapidly rotating Be stars, in which equatorial confined non-radiial modes are usually acting. So far, for one Be star, namely \(\mu\) Cen, a possible relation between mass-ejection and the beating of several non-radiational pulsation modes has been found (Rivinius et al. 1998).

It is also interesting to note that the mode splitting due to the rotation is clearly present in the pulsation properties, although a 6 d period is favored by their analysis (Telting et al. 1997). It would be worth examining whether the presence of magnetic field can also be traced back in the pulsation modes. If so, this will give a strong constraint on the evolutionary status of \(\beta\) Cep.

Several other issues are still to be solved. First of all, why this star has a magnetic field cannot be answered: \(\beta\) Cep does not belong to the helium-peculiar stars (Rachkovskaya 1990), which are known to have strong magnetic fields, see e.g. Bohlender et al. (1987).

Second, the origin of the variable \(\text{H}\alpha\) emission is not known, and it is worthwhile to investigate a possible correlation with the magnetic field behavior. Whether the much higher value of the average magnetic field value obtained of \(R\) & Kemp (1978) in November 1976 was significant cannot be checked, although it should be noted that this value is still consistent with noise, considering that 7 of their individual measurements are within 1\(\sigma\), 2 within 2\(\sigma\) and 1 within 3\(\sigma\). At any rate, at this epoch there are no records of emission in the \(\text{H}\alpha\) line. The magnetic measurements presented by Henrichs et al. (1993) were taken during a clear emission phase. The error bars of 150 \(\text{G}\) of these and consequently taken observations by Hill (private communication) are in retrospect too large too have revealed a field strength compatible with the values we report in this paper. This explains why no 12 d periodicity could be found in these magnetic data. Continuing long-term monitoring of the magnetic field is obviously needed to see whether there is any correlation with the \(\text{H}\alpha\) behavior.

Third, we have fitted a simple sine curve through the magnetic data. This is obviously a first approximation, and when more measurements become available, a search for deviations of a sine curve, as is found for most magnetic stars, can be done.

Fourth, the preliminary modeling in Section 6.6 implies a 12 day modulation in the X-ray flux. A quantitative study is needed to confirm this prediction.
Last, it should be noted that the star $\beta$ Cep appears to be one of the very few stars in its class which shows this type of strong wind variability and in this respect $\beta$ Cep is an exceptional $\beta$ Cep star.

Acknowledgements. We thank John Telting and Gautier Mathys for their constructive comments and discussion. The helpful assistance of the observatory staff members at TBL, OHP, GSFCC and Vilspa is well remembered and greatly acknowledged. We thank Rens Waters for his helpful comments. JDJ acknowledges support from the Netherlands Foundation for Research in Astronomy (NFRA) with financial aid from the Netherlands Organization for Scientific Research (NWO) under project 781-71-053. GAW acknowledges support from the Natural Sciences and Engineering Council of Canada (NSERC) in the form of an NSERC postdoctoral fellowship held during the course of this work.

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Chapter 7

Modeling cyclical Hα line profile variability in O-type stars

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To be submitted to A&A

Abstract. The observed cyclical wind variability in O-type stars is thought to be due to the presence of Corotating Interaction Regions (CIRs) in the stellar wind. In this paper we investigate the impact of CIR-like wind structures on the formation and shape of the Hα line in O-star spectra. We use a semi-empirical 2D radiative transfer model in which line profiles are synthesized by computing for each velocity bin the flux from a constant line-of-sight velocity. The wind structure is defined by one or more spiral structures, which have parameters for the curvature, width, length, and radius of maximum perturbation. A dynamical spectrum is simulated by computing a series of model spectra in which the structure is rotated around the star.

The goal of this exercise is to find out whether the model spectra can reproduce the variations observed in the Hα line for different O stars. If so, in principle one could invert the problem and derive the wind structure from the Hα line variations. It turns out that our model spectra are capable to qualitatively explain the typical variations encountered in the Hα profiles. Although we managed to reconstruct the input wind structure for a set of test data with help of a genetic algorithm fitting procedure, it was not possible to successfully apply this method to a set of observed spectra.

7.1. Introduction

Massive stars (spectral types O and B) have strong stellar winds driven by radiation pressure. Careful investigations have shown that nearly all O and many B stars exhibit cyclical line-profile variations in their ultraviolet and visible spectrum which are due to variations in their stellar winds (many other classes of stars also show cyclical wind variations; for a recent overview, see Kaper & Fullerton 1998). The most prominent type of variability in OB-star winds is caused by the accelerating discrete absorption components (DACs) in ultraviolet resonance lines. These DACs appear as broad absorption features at low or intermediate (0.2-0.5v∞) line-of-sight velocity and narrow in width when they approach the terminal velocity. We refer to Kaper et al. (1996, 1999a) for an extensive discussion on the DAC behavior in 10 different O stars. A key observation is that DACs appear periodically (Henrichs et al. 1988, Prija 1988), and that their recurrence timescale is shorter in stars with a higher (projected) rotational velocity. This strongly suggests that stellar rotation determines the timescale of variability in OB-star winds.

Similar cyclical variability is observed in the Hα line of many O-type stars (Ebbets 1982, Kaper et al. 1998, 2000). In practice only (super)giants produce observable wind variability in the Hα line; the higher luminosities of these stars result in higher mass-loss rates. The variations in the Hα profile are often much more complex than those encountered in the UV resonance lines. Changes occur both in absorption and in emission, and various features are present that move back- and/or forwards in velocity. Other wind-or transition region lines like He I λ5876 and He II λ4686 show a similar behaviour (e.g., Leep & Conti 1979, Henrichs 1991, Reid & Howarth 1996, Fullerton et al. 1991, Henrichs et al. 1998b, Chapter 3).

Only in a small number of cases, the ultraviolet and optical spectrum has been observed simultaneously in order to investigate whether wind variations close to the stellar surface (as traced by e.g. the Hα line) and further-out in the wind (derived from the UV resonance lines) are physically connected. Kaper et al. (1997) performed coordinated Hα and ultraviolet spectroscopy of the brightest northern-hemisphere O-stars. They showed that the variations observed in the Hα line are linked to the DAC behaviour in UV resonance lines. A more extensive multi-site and multi-wavelength campaign was conducted in October 1994 on the O7.5 III star ζ Per (de Jong et al., 2000a, Chapter 2). The obtained time series demonstrate that the wind variability is strictly periodic at 2.09 days in all studied lines. The results of this campaign support the conclusion that we are witnessing the same phenomenon in the optical and ultraviolet wind lines. Due to the switch-off of the International Ultraviolet Explorer (IUE), in November 1996 (MUSICOS, Henrichs et al. 1998b, Chapter 3) only the optical line variations could be studied.

The periodical behavior of the variability in O-star winds on a timescale that can be identified with the stellar rotation period strongly suggests that one or more large-scale structures in the stellar wind are in corotation with the
star. These structures must be stable for at least several rotation periods. Currently, the best explanation for the cyclical wind variability is given by the Corotating Interaction Regions (CIR) model in which the emergent wind flow is perturbed by some structure at the surface of the star. Such a structure causes a local change in the radial flow properties, for example a higher density and a smaller outflow velocity. Due to the stellar rotation the wind streams are curved. Further-out in the wind, the slow (or fast) moving stream interacts with the ambient flow resulting in a shock front that corotates with the star. This model was first proposed by Mullan (1984) in analogy with the CIRs in the solar wind. Cranmer & Owocki (1996) made the first hydrodynamic computations based on this model. The resulting variations in the calculated resonance line profiles indeed closely resemble the observed DAC behavior. They found that the primary contribution to the (Sobolev) optical depth comes from a region trailing the CIR, where a strong gradient in the wind velocity (“kink”) is present. In this model the origin of the perturbation at the stellar surface (a bright spot giving rise to a local increase in wind mass-loss rate) is not specified, but could be caused by e.g. the presence of a surface magnetic field and/or non-radial pulsations (NRPs).

The hydrodynamical model by Cranmer & Owocki (1996) is able to reproduce the DAC behavior in UV resonance lines. However, they did not attempt to model the H\textalpha line. Although H\textalpha shows a much larger complexity in its variability pattern, and is more difficult to model than the UV resonance lines, it probes the part of the stellar wind where the seed of the variability has to be found. Another advantage of using the H\textalpha line as a diagnostic for the detailed structure of OB-star winds is that this line is accessible from the ground and that with present-day instrumentation H\textalpha spectra can be obtained of massive stars in external galaxies up to the Virgo cluster (Kudritzki 1998). Furthermore, for stars with dense stellar winds the UV resonance lines are saturated which limits the study of wind variability to lines like H\textalpha.

To study the impact of corotating structures (like those corresponding to CIRs) on the formation and variability of the H\textalpha line, we developed a simple semi-empirical model. With this model we compute time series of spectra for a large variety of wind structures. We investigate which structures can reproduce the different features observed in the line spectra. Since H\textalpha is formed close to the star, stellar rotation has to be taken into account. The fact that H\textalpha is a subordinate line where the populations are dominated by recombination, the line source function is proportional to the square of the wind density (in stead of proportional to the wind density in the case of a resonance line). Therefore, the contribution of wind emission to the H\textalpha line is relatively much more important compared to the case of resonance-line formation. For the same reason, wind material outside the line of sight towards the star contributes to the H\textalpha line. Due to the projection of the wind velocity photons originating at different locations in the stellar wind (in- and outside the line of sight) can contribute to a given position in the H\textalpha line. In UV resonance lines wind variability is limited to the blue-shifted absorption part of the profile (and thus to material contained in the line of sight where projection effects are negligible).

In section 7.2 the semi-empirical model is described which we used to produce model H\textalpha spectra. In section 7.3 some model spectra are compared to H\textalpha spectra which were obtained for some giant and supergiant O stars. The predicted time evolution of the H\textalpha line for a given large-scale structure in the stellar wind is shown in section 7.3.2. We demonstrate that the (CIR) model has the potential to qualitatively explain the observed H\textalpha-line variability (Sect. 7.4). We tried to invert the problem and derive the wind structure by modelling the observed time series of spectra using a genetic algorithm fitting procedure (Sect. 7.5). Although we managed to reconstruct the wind structure used to produce an artificial set of spectra, an application of this method to an observed set of spectra was not successful. In the last section we summarize our conclusions and discuss possible future improvements of our model.

7.2. Description of the model

7.2.1. Outline

The line profiles are computed by solving in each velocity bin the loci of constant line-of-sight velocity and calculating the out-coming flux using the radiative transfer equation in the Sobolev approximation. For this we use a \beta velocity law for the radial velocity and conservation of angular momentum for the tangential velocity component. The density is computed using the continuity equation and a phase dependent perturbation. For the source function and optical depth, we assume appropriate relations for the absorption and emission coefficients (\propto \rho^2 for a recombination line). The photospheric intensities are approximated by Doppler-shifted Gaussian line profiles.

The density structure is parameterized by means of a 2-dimensional grid \Delta \rho(r, \phi) where r and \phi are respectively the distance to the stellar center and angle in the plane of rotation; \Delta \rho is the perturbation in density relative to the smooth “underlying” wind. We do not take into account any changes in the velocity field due to these perturbations. Although this violates the continuity equation we assume that the velocity changes that would be induced otherwise are small enough to ignore.

7.2.2. Line formation

This section describes the computation of a line profile given the wind and radiative transfer parameters. Here, we consider the line formation only in a 2-dimensional plane through the equator of the star which is assumed to be seen edge-on. The standard Cartesian coordinate system with impact parameter p and the line of sight z is used (e.g.
Lamers et al. 1989), as well as a polar coordinate system
\((r, \phi)\) for the radially defined quantities \(r = \sqrt{p^2 + z^2}\)
and \(\tan \phi = z/p\). All distances are in units of stellar radii
\((R_\star)\) (see Fig. 7.1).

**Fig. 7.1.** Coordinate system used in our model. The \(p, z\) and \(r\)
coordinates are measured in stellar radii \((R_\star)\) and \(\phi\) is in radians.
Furthermore, some contours of equal line-of-sight velocity are
drawn (the dashed lines indicate negative line-of-sight velocity).
The direction of rotation is clockwise and the observer looks
from below.

For each bin \(i\) in the line profile, the corresponding
radial velocity \(v_{\text{obs}}\) \((\text{km} \, \text{s}^{-1})\) and flux \(F_i\) are calculated.
We first compute for a given number of impact rays \(p_j\) for
each velocity \(v_{\text{obs}}\) the loci of constant velocity \(z_{ij}\) by solving
the equation describing the 2-dimensional velocity field
including rotation:

\[
v_{\text{obs}}(p, z) = v_r(r) \frac{z}{r} - \left( \frac{v_{\text{rot}}}{r} \right) \left( \frac{p}{r} \right).
\]  

(7.1)

Here \(v_{\text{rot}}\) is the projected rotational velocity and \(v_r(r)\) the
radial wind velocity which is determined by the \(\beta\) velocity
law (e.g. Puls et al. 1996):

\[
v_r(r) = v_\infty \left( 1 - \frac{b}{r} \right)^{\beta}.
\]  

(7.2)

\[
b = 1 - \left( \frac{v_0}{v_\infty} \right)^{1/\beta},
\]  

(7.3)

where \(v_0\) is the velocity at the base of the wind. For this
we take the sound speed \(v_{\text{sound}} = \sqrt{kT/\mu m_H}\); \(v_\infty\) is the
terminal velocity of the wind. An example of the adopted
2-dimensional velocity field is demonstrated in Fig. 7.2.

Subsequently, a matrix \(z_{ij}\) is calculated containing the
\(z\) coordinate for \(N_j\) velocity bins \((v_{\text{obs}})\) in the line profile
and \(N_j\) impact rays \(p_j\). For each point \((p_j, z_{ij})\) the density
\(\rho\) is computed:

\[
\rho = \frac{v_0 \Delta \rho(r, \phi)}{r^2 v_r(r)}.
\]  

(7.4)

Here \(v_0\) is the same constant as in Eq. 7.3 which means that
the unperturbed density at \(r = 1\) is set to 1. \(\Delta \rho(r, \phi)\) is the
density perturbation (see Sect. 7.2.3) and is equal to 1 in
case of no perturbation. It is the only \(\phi\) dependent quantity
and accounts for all the variability. The optical depth \(\tau\) in
the Sobolev approximation is given by:

\[
\tau = \kappa \rho^n \left| \frac{\partial v_{\text{obs}}}{\partial z} \right|^{-1},
\]  

(7.5)

where \(n\) depends on the physics of the radiative transfer in
the line concerned; \(n = 2\) for the recombination line \(\text{H}\alpha\)
and the subordinate UV line of \(\text{N}\,\text{I}\); \(n = 1\) for the UV
resonance lines. \(\partial v_{\text{obs}}/\partial z\) is computed from Eq. 7.1. The intensity \(I\) is obtained by solving the equation of radiative transfer:

\[
I = S (1 - e^{-\tau}) + I_{\text{phot}} e^{-\tau}
\]  

(7.6)

\[
I_{\text{phot}} = \left\{ \begin{array}{ll}
1 - G(p, v_{\text{obs}}) \cdot L(p) & \text{for } p \leq 1 \\
0 & \text{for } p > 1
\end{array} \right.
\]  

(7.7)

\[
G(p, v_{\text{obs}}) = I_{\text{min}} \exp \left[ - \left( \frac{v_{\text{rot}}(p) - v_{\text{obs}}}{v_{\text{th}}} \right)^2 \right]
\]  

(7.8)
\[ L(p) = \frac{3}{5} \left( \sqrt{1 - p^2} + \frac{2}{3} \right). \tag{7.9} \]

Here \( I_{\text{phot}} \) is the photospheric intensity which is assumed to have a Gaussian-shaped line profile \( (G(p, v_{\text{obs}})) \) with a width \( v_{\text{obs}} \). This profile is Doppler-shifted by the rotational velocity at the impact parameter \( p(v_{\text{rot}}(p)) \). Furthermore, \( I_{\text{phot}} \) is reduced by a grey-atmosphere limb-darkening factor \( L(p) \). Both \( I_{\text{phot}} \) and \( S \) are taken as a fraction of the continuum intensity \( I_c \), which is not specified in this paper.

We assume that the absorption coefficient \( \kappa \) and the source function \( S \) are constant. This is a reasonable assumption for \( \text{H}\alpha \), since Puls et al. (1996) showed from numerical computations that the departure coefficients of \( \text{H}\alpha \) only weakly depend on \( r \) and \( v_r(r) \). We hereby also assume that the wind temperature is constant. The constants \( \kappa, S \) and \( I_{\text{min}} \) in these equations are determined empirically by fitting a line profile to the average observed profile (see Sect. 7.3.1).

We further assume that the perturbations do not affect the velocity field (Eq. 7.1). Introducing a density-velocity relation in accordance with the continuity equation would require a hydrodynamical approach, which is beyond the scope of this paper. In Sect. 7.6 we discuss this further. Furthermore, the radiative transfer equation (Eq. 7.6) is only evaluated along the lines of constant \( v_{\text{obs}} \) assuming the validity of the Sobolev approximation everywhere. Finally, the flux for each velocity bin \( i \) is computed by integrating \( I \) numerically (using the extended trapezium rule) along the lines of constant \( v_{\text{obs}} \):

\[ F(v_{\text{obs}}) = \int_{P_{\text{min}}}^{P_{\text{max}}} I(v = v_{\text{obs}}) dp = \frac{1}{2} I_{1,1} + \frac{1}{2} I_{1,3} + \sum_{i=2}^{N_{\text{bin}}} I_{i,k} \tag{7.10} \]

### 7.2.3. Density perturbation

The density perturbation is assumed to be time-invariant in the corotating frame of the star. All the variations are due to the rotation of the wind structure in the observer's frame. The transformation from the corotating frame (polar coordinates) to the observers frame (Cartesian coordinates) is given by:

\[ \phi = \frac{1}{2\pi} \arctan \left( \frac{z}{p} \right) + \Phi + \frac{1}{4} \]
\[ r = \sqrt{p^2 + z^2}. \tag{7.11} \]

In this equation both \( \phi \) and \( \Phi \) run from 0 to 1. \( \Phi \) is the rotational phase of the structure, which is equal to the \( \phi \) coordinate of the structure at the line \( (p = 0, z < 0) \). If the stellar rotation period \( P \) is known the rotational phase is defined as:

\[ \Phi = \frac{t}{P} - \text{INT} \left( \frac{t}{P} \right), \tag{7.12} \]

where the function \( \text{INT}(x) \) returns the largest integer value less than \( x \). We define the perturbation as consisting of \( N_{\text{blob}} \) blobs in \((r, \phi)\) coordinates:

\[ \Delta \rho(r, \phi) = 1 + \sum_{i=1}^{N_{\text{blob}}} (\Delta \rho_{\text{max}}) \times P(r, \phi) \tag{7.13} \]

\[ P(r, \phi) = \exp \left\{ \left( \frac{r - r_{01}}{\sigma_r} \right)^2 - \left( \frac{\Delta \phi}{\sigma_\phi} \right)^2 \right\} \tag{7.14} \]

\[ \Delta \phi = \phi - (\phi_{01} + (r - 1)\phi_1) \tag{7.15} \]

The density distribution \( P(r, \phi) \) of a blob is described by a Gaussian distribution function: \( \Delta \rho_{\text{max}} \) is the maximum amplitude of a blob, \( \sigma_r \) and \( \sigma_\phi \) are the length and width, \((r_0, \phi_0)\) is the position of the density maximum. Finally, \( \phi_1 \) determines the deflection in \( \phi \), which transforms into a curvature in the Cartesian \((p, z)\) coordinates. So, these blobs form Archimedian spiral structures in the wind, resembling CIRs.

The above equation can use up a considerable amount of computing time when exploited in the inverted problem (section 7.5). In order to save this computing time we computed the perturbation only on a fixed \( 50 \times 50 \) grid. The value for a certain \((p, z)\) point is obtained by linear interpolation using the three closest points. The grid is large enough to ensure that interpolation errors are negligible.

To save even more computing time most coefficients used in these interpolations are computed beforehand for each \((p, z, \Phi)\) point. Finally, this perturbation is put into Eq. 7.4.

### 7.3. Model spectra

In this section the wind parameters are empirically derived by fitting the average line profiles of \( \xi \) Per (an O7.5 giant) and \( \alpha \) Cam (an O9.5 supergiant). The impact of the adopted large-scale structure on the detailed shape of the \( \text{H}\alpha \) profile is investigated by computing models for a large range of structure parameters.

#### 7.3.1. General wind parameters

First, appropriate values are chosen for \( v_0, v_\infty, \) and \( \beta \) for the star in question. Non-perturbed models are compared to the average \( \text{H}\alpha \) spectrum of a time-series, which is assumed to be close to the unperturbed profile, in order to empirically derive values for \( S, \kappa \) and \( I_{\text{min}} \).

We have done this for two stars: \( \xi \) Per (O7.5III) and \( \alpha \) Cam (O9.5Ia) as examples of the occurrence of wind variability in the \( \text{H}\alpha \) profile of giants and supergiants, respectively. For \( \xi \) Per we used the line profiles obtained during the October 1994 campaign (De Jong et al. 2000a, Chapter 2, Fig. 7.3). The observations of \( \alpha \) Cam were taken in October 1995 with the 1.52m telescope at the Observatoire de Haute Provence (OHP) and with the JKT.
at the Roque de los Muchachos observatory on La Palma (Kaper et al. 2000).

In the velocity law, we use $\beta = 1$ which is representative of the values found by Puls et al. (1996) and Petrenz & Puls (1996) for different O stars. For $\xi$ Per we used $v_\infty = 2330$ km s$^{-1}$ (Kaper et al. 1999a), $v_{\text{rot}} = 200$ km s$^{-1}$ (Penny 1996) and 25 km s$^{-1}$ for both $v_0$ and $v_{\text{th}}$, which is the sound speed at $T = 36000$ K. We thus assume that the profile is also broadened by some turbulence. By comparing several models with different $S$ and $\kappa$ to a good template of the October 1994 data of this star we could empirically determine their value. Fig. 7.4 shows some models for different values of $\kappa$ and $S = 0.2$. The best fit to the average profile is obtained for $\kappa = 1.7 \times 10^4$. Increasing $S$ is not a good option, because that would make the profile much narrower and increase the red wing emission. This is illustrated in Fig. 7.4 by the dash-dotted profile, which is computed for $S = 0.5$ and $\kappa = 6 \times 10^4$. Since the line profile is visible up to about 600 km s$^{-1}$, H$\alpha$ must be formed up to 1.4 $R_\odot$.

The H$\alpha$ line of $\alpha$ Cam is predominantly in emission. The UV resonance lines are completely saturated which leaves strong subordinate lines like H$\alpha$ lines the only alternative to study the wind variability. This star has $v_\infty = 1590$ km s$^{-1}$ (Kaper et al., 1999a) and $v_{\text{rot}} = 129$ km s$^{-1}$ (Howarth et al., 1997). The line profile (see Fig. 7.5) shows that the wind emission extends up to 1000 km s$^{-1}$, which means (assuming $\beta = 1$) that H$\alpha$ is formed up to 2.7 $R_\odot$. A reasonable fit was achieved for $\kappa = 3 \times 10^5$ and $S = 0.7$ (see Fig. 7.5). With lower values for $S$ the profile could not be fitted, because the blue shifted absorption in the P-Cygni profile becomes too large. Note that in our model $\kappa$ contains both the density at the base of the wind and the absorption coefficient. E.g., differences in mass-loss rate between $\xi$ Per and $\alpha$ Cam do not come out straightforwardly. But deriving the (average) wind parameters from a fit to the H$\alpha$ line is not the scope of this paper.

7.3.2. Characteristic variations

In order to get an idea of the H$\alpha$ line profile variability due to a given large-scale structure in the stellar wind, we ran test models and produced dynamic spectra for 120 different structures by varying the parameters defined in section 7.2.3. A dynamic spectrum consists of model spectra which are stacked on top of each other as a function of rotational phase. To enhance the contrast, the model spectra are divided by the model spectrum without perturbation. The (residual) flux is represented in levels of gray: black means additional absorption or less emission, white additional emission or less absorption. The explored parameter space is given by: $\Delta \rho_{\text{max}} = 2$ for the maximum increase in density, $\sigma_r \in \{0.05, 0.1, 0.2\}$ and $\sigma_\phi \in \{0.05, 0.15\}$ for the length, respectively width of the blob, $r_0 \in \{1.0, 1.1, 1.2, 1.3, 1.4\}$ for the position of the blob (distance of density maximum above the stellar surface plus
Table 7.1. Structure parameters for the models in Fig. 7.6.

<table>
<thead>
<tr>
<th>model</th>
<th>$\Delta \rho_{\text{max}}$</th>
<th>$\phi_0$</th>
<th>$\phi_1$</th>
<th>$\sigma_\phi$</th>
<th>$r_0$</th>
<th>$\sigma_r$</th>
</tr>
</thead>
<tbody>
<tr>
<td>(a)</td>
<td>2.00</td>
<td>0.50</td>
<td>0.00</td>
<td>0.03</td>
<td>1.00</td>
<td>0.10</td>
</tr>
<tr>
<td>(b)</td>
<td>2.00</td>
<td>0.50</td>
<td>0.00</td>
<td>0.03</td>
<td>1.20</td>
<td>0.10</td>
</tr>
<tr>
<td>(c)</td>
<td>2.00</td>
<td>0.50</td>
<td>0.00</td>
<td>0.03</td>
<td>1.30</td>
<td>0.10</td>
</tr>
<tr>
<td>(d)</td>
<td>2.00</td>
<td>0.50</td>
<td>0.50</td>
<td>0.03</td>
<td>1.30</td>
<td>0.10</td>
</tr>
<tr>
<td>(e)</td>
<td>2.00</td>
<td>1.00</td>
<td>0.00</td>
<td>0.03</td>
<td>1.30</td>
<td>0.10</td>
</tr>
<tr>
<td>(f)</td>
<td>2.00</td>
<td>1.00</td>
<td>1.00</td>
<td>0.03</td>
<td>1.20</td>
<td>0.10</td>
</tr>
</tbody>
</table>

1), and $\phi_1 \in \{0.0,0.5,1.0,2.0\}$ for the curvature of the structure. We also changed the value of $\beta$ in the velocity law and of the stellar rotation velocity for a few models; $v_{\infty}$ was kept fixed at the value for $\xi$ Per.

Fig. 7.6 presents an overview of the most characteristic features encountered in the dynamical spectra based on these models. The parameters of the selected models are given in Table 7.1. Most of the dynamic spectra show features which move from blue to red due to rotation (see Fig. 7.6(a) for an example). These variations originate very close to the stellar surface where the radial velocity due to wind expansion is relatively low. In order to get features which move from red to blue due to the expanding wind (as is observed in several cases), the point where the density of the blob reaches a maximum must be further out in the wind: we found that $r_0 \geq 1 + \sigma_\tau$, i.e. the density maximum must be located at a distance at least $\sigma_\tau$ above the stellar surface. In almost all cases a blue-ward moving feature is still accompanied by a red-ward moving feature (induced by the stellar rotation). To get the variations in the dynamic spectrum to appear only in the region dominated by wind expansion, $r_0$ must be greater than $3\sigma_\tau + 1$. These features are either bow-like (see Fig. 7.6(c)) or only blue-ward moving (see Fig. 7.6(d) and (e)) if $\phi_1 \geq 0.5$.

In order to investigate the effect of changing $\beta$ and $v_{\text{rot}}$ we computed the models from Fig. 7.6(b) and (e) also for $\beta = 1.3$ and $v_{\text{rot}} = 20 \text{ km s}^{-1}$ (instead of 1 and 200 km s$^{-1}$, respectively). The results are shown in Fig. 7.7. One can see that changing $\beta$ affects the velocity of the feature and its "acceleration". In principle, this could also be reached by selecting another value for $r_0$ and $\phi_1$. Therefore, we avoid ambiguity by fixing $\beta$ to the known values based on other studies. Changing $v_{\text{rot}}$ has only effect on features which are formed close enough to star (lower panels in Fig. 7.7), i.e. $r_0 \lesssim 1 + \sigma_\tau$.

Finally, we made a more complicated model with three blobs (CIRs) for which the parameters are given in Table 7.2. This model is shown in Fig. 7.8. In case of a low value for the source function $S = 0.2$ the different CIRs are easily recognizable as separate components in the spectrum. In Sect. 7.3.1 we derived this value for $\xi$ Per which has an H$\alpha$ profile mainly in absorption. However, the dynamic spectrum becomes much more complicated if $S$ is higher. In that case the emission of one component starts to fill in the absorption of another. This results into the complicated spectrum in the right panel of Fig. 7.8 for which $S = 0.5$.

Table 7.2. Configuration of the multiple CIR test model. In total three CIRs were used with the following parameters.

<table>
<thead>
<tr>
<th>No.</th>
<th>$\Delta \rho_{\text{max}}$</th>
<th>$\phi_0$</th>
<th>$\phi_1$</th>
<th>$\sigma_\phi$</th>
<th>$r_0$</th>
<th>$\sigma_r$</th>
</tr>
</thead>
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<tr>
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</tr>
<tr>
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<td>3</td>
<td>0.75</td>
<td>0.5</td>
<td>0.05</td>
<td>1.3</td>
<td>0.2</td>
</tr>
</tbody>
</table>

7.4. Modeling the H$\alpha$ variations in $\xi$ Per and $\alpha$ Cam

7.4.1. $\xi$ Per (07.5III)

A unique set of almost continuous H$\alpha$ observations of the 07.5 III star $\xi$ Persei (HD 24912) over a period of 10 days was obtained in October 1994 (de Jong et al., 2000a, Chapter 2). The period of the cyclical variations is 2.09 days. However, the interpretation of the DAC behavior in UV resonance lines suggests that the rotation period of $\xi$ Per is twice as long, namely 4.18 days (de Jong et al. 2000a, Chapter 2). In Fig. 7.10 the observed dynamic H$\alpha$ spectrum (consisting of 314 spectra) is shown folded with the 4.18 d period using Eq. 7.12.

We estimated the shape, size and location of the blobs based on a qualitative comparison between different sets of
Fig. 7.6. These 6 models give a summary of the characteristic features which we encountered in the 120 different models. Model (a) is computed with $\Delta \rho_{\text{max}} = 2$, $\phi_0 = 0.5$, $\phi_1 = 0.0$, $\sigma_\phi = 0.05$, $r_0 = 1.0$ and $\sigma_r = 0.1$. The upper panel shows the line profiles, the middle panel shows a dynamic spectrum containing the line profiles divided by the non-perturbed profile (relative flux in levels of grey). The shape of the structure (at phase 0.75, rotation is clockwise, observer looks from below) is shown in the lowest panel (in cartesian coordinates). It is an elongated blob which starts at the surface. In model (b) ($r_0 = 1.2$) and (c) ($r_0 = 1.3$) the blob is positioned further away from the star, resulting in “bow” shaped features in the dynamic spectrum. Models (d), (e) and (f) are curved ($\phi_1 = 1.0$) and yield blue-ward moving features in the dynamic spectrum.
model spectra with the observed ones. We adopted $\beta = 1$, $S = 0.5$, and the stellar parameters listed in 7.3.1. Although $S = 0.2$ followed from the fit of the average line profile (see Sect. 7.3.1), this value turned out to be too low to explain the amount of relative emission in the residual spectra. Fig. 7.9 depicts the location and shape of the perturbations that result in a dynamic spectrum qualitatively similar to the observations (see Table 7.3 for the chosen parameters). The strong absorption around $v = -300$ km s$^{-1}$ and $\Phi \simeq 0.4$ is caused by two "CIRs" centered at $r_0 = 1.2R_\star$, with a width $\sigma_r = 0.03$, a length $\sigma_\phi = 0.1$ and a curvature $\phi_1 = 0.8$. These CIRs are also responsible for the emission around $\Phi \simeq 0.4$ and $\Phi \simeq 0.7$ when they reside next to star. The red-shifted absorption feature that appears at $\Phi \simeq 0.4$ is much more difficult to explain. It could be caused by a blob close to the star, the corresponding red-ward moving absorption feature of which coincides with the emission from another blob. In that case the emission will cancel the absorption on the red side, leaving only absorption on the blue side. Therefore, we put two blobs near the surface in between the strong CIRs, but even these cannot explain the whole picture. Besides, the found values for $\phi_1$ are much larger than the curvature found from hydrodynamical computations (Cramer & Owocki 1996). We will discuss this further in Sect. 7.6.

![Fig. 7.8. Model containing multiple CIRs for $S = 0.2$ (left panel) and $S = 0.5$ (right panel).](image)

Table 7.3. Structure parameters for model of $\xi$ Per.

<table>
<thead>
<tr>
<th>No.</th>
<th>$\Delta \rho_{\text{max}}$</th>
<th>$\phi_0$</th>
<th>$\phi_1$</th>
<th>$\sigma_\phi$</th>
<th>$r_0$</th>
<th>$\sigma_r$</th>
</tr>
</thead>
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<td>0.8</td>
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<td>0.1</td>
</tr>
<tr>
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<td>0.03</td>
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</tr>
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<td>0.1</td>
</tr>
<tr>
<td>4</td>
<td>1.5</td>
<td>0.1</td>
<td>0.3</td>
<td>0.03</td>
<td>1.0</td>
<td>0.1</td>
</tr>
</tbody>
</table>

7.4.2. $\alpha$ Cam (O9.51a) and $\zeta$ Ori (O9.71b)

The supergiants $\alpha$ Cam (HD 30614) and $\zeta$ Ori (HD 37742) were not observed during a campaign with continuous coverage; a number of Hα spectra covering a period of 8 days were obtained spectra each night (Kaper et al. 2000). The night averages of these spectra are plotted in Fig. 7.11. Both stars vary on a timescale of several days; these timescales are of the same order as the total length of the time series (8 nights) so that accurate periods cannot be derived. In the case of $\alpha$ Cam a CLEANed Fourier Transform shows a strong peak at a period of $5.6(5)$ d (see Kaper et al. 2000); several other peaks appear in the power diagram as well. For $\zeta$ Ori Kaper et al. (2000) find a peak in the power diagram at $7.1(8)$ d. In both stars the variations are qualitatively similar in character and are concentrated in the range $-500$ and $500$ km s$^{-1}$. The blue-shifted absorption and emission
Modeling cyclical Hα line profile variability in O-type stars

**Fig. 7.10.** Left panel: Dynamic spectrum of ξ Persei (October 1994) folded with a period of 4.18 days. Right panel: Model dynamic spectrum with estimated structure.

**Fig. 7.9.** Estimated wind structure for ξ Per based on a comparison of observed and modeled Hα spectra.

in the P-Cygni-type profiles show dramatic variability. In ζ Ori the main emission peak varies strongly, while sometimes a blue-shifted emission peak appears superposed on the P-Cygni absorption.

We calculated model spectra using the wind parameters derived in Sect. 7.3.1 and two structures with a different curvature (see Table 7.4). The line profiles of the model are shown in the right panel of Fig. 7.11 (a 6-day period was taken for comparison). It turns out that close to the star τ is very large and Eq. 7.6 becomes \( I \approx S \) (i.e. \( I_{\text{phot}}e^{-T} \) gets very small). Therefore, all the variability has to be due to variations in wind emission alone. In this way the moderate emission variability in α Cam can be explained quite well. But in ζ Ori the Hα variations are so large that the emission at \( t = 14 \) almost completely disappears. This is difficult to explain with just a reduction in wind emission.

**Table 7.4.** Structure parameters for model of α Cam.

<table>
<thead>
<tr>
<th>No.</th>
<th>( \Delta p_{\text{max}} )</th>
<th>( \phi_0 )</th>
<th>( \phi_1 )</th>
<th>( \sigma_\phi )</th>
<th>( r_0 )</th>
<th>( \sigma_r )</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
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<td>0.05</td>
<td>1.2</td>
<td>0.2</td>
</tr>
<tr>
<td>2</td>
<td>5</td>
<td>0.3</td>
<td>0.2</td>
<td>0.05</td>
<td>1.2</td>
<td>0.2</td>
</tr>
</tbody>
</table>
Fig. 7.11. Time series of Hα observations of α Cam and ζ Ori obtained in October 1995. For comparison, we show one series of model spectra which qualitatively show a similar, though not identical, behaviour.

7.5. Inferring the wind structure from fitting dynamic spectra

We investigated the possibility of deriving the wind structure from an observed dynamic spectrum. The dynamic spectrum containing $F(v_{\text{obs}}, t)$ is folded with the rotational period $P$ using Eq. 7.12 which results in $F(v_{\text{obs}}, \Phi)$. The spectra are rebinned to equally spaced phases using a triangular weight function. A matrix $F_{ij}$ is obtained containing the fluxes for $N_v$ velocity- and $N_j$ phase bins. Finally, the individual spectra are divided by an appropriate template profile, which should represent the unperturbed line profile.

For each phase bin $j$ a model is computed as described in Sect. 7.2 from which a model dynamic spectrum $M_{ij}$ is constructed. In order to test how good the model fits the observations we first compute the squared difference matrix:

$$D_{ij} = \left( \frac{F_{ij} - M_{ij}}{\sigma_{ij}} \right)^2$$  \hspace{1cm} (7.16)

Here $\sigma_{ij}$ denotes the noise in the observed flux. We evaluate the average $<D>$ and standard deviation $\sigma_D$ of this matrix and require that the best fit has the highest value for

$$f = \frac{1}{<D>} + \zeta \sigma_D$$  \hspace{1cm} (7.17)

Here $\zeta$ is a user adjustable parameter, which determines the importance of the $\sigma_D$ criterium. It is quite likely that for each dynamic spectrum a unique structure can be found, since both quantities depend on two variables and each line-of-sight velocity can be traced back to a single curve in the wind.

We use the Genetic Algorithm (GA) technique (Charbonneau (1995), Wall (1995) or Goldberg (1989)) to fit the observations. This is a general purpose optimization algorithm which uses principles from natural evolution. The algorithm was first developed by Holland (1975) and the robustness of the algorithm has been mathematically proven (Goldberg 1989 for a good review). We used the so-called Steady State GA which works as follows:

1. All the parameters in Eq. 7.13 are stored in a single string of bits which is called a genome. Each parameter occupies $n_{\text{bit}}$ bits of this string which is translated to a value between $p_{\text{min}}$ and $p_{\text{max}}$ as follows:

$$p = p_{\text{min}} + \frac{p_{\text{max}} - p_{\text{min}}}{2^n - 1} \sum_{q=1}^{n} y_q 2^q$$  \hspace{1cm} (7.18)

Where $y_q$ is the value of a bit (0 or 1). So, the genome contains in total $6N_{\text{blob}} \cdot n_{\text{bit}}$ bits. See Fig. 7.12 for an example. For $n_{\text{bit}}$ we took 6, so that each parameter can obtain one of 64 values (this turned out to be sufficient). The values for $p_{\text{min}}$ and $p_{\text{max}}$ have to be confined to a sensible range for each of the 6 parameters describing a blob, hereby carefully considering the observations which are to be fitted. Furthermore, an "individual" is defined as an object which contains both a filled genome (the genotype) and the resulting model (the phenotype).

2. A population of $N_{\text{pop}}$ individuals are created by filling their genomes randomly and computing the resulting models (the phenotype). These models are compared with the observations and assigned a fitness $(f)$, which is the result of Eq. 7.17.

3. Individuals are selected for reproduction according to their fitness. We use for this the Stochastic Remainder selection scheme (Goldberg, 1989). In this scheme all individuals are assigned a probability for selection (fitness): $p_i = f_i / \sum f_i$. Then the expected amount of selections $e_i = N_{\text{pop}} p_i$ is computed. A temporary population is created for which individual $i$ is selected at least INT$(e_i)$ times and possibly once more with a probability defined by the fractional part of $e_i$. $N_{\text{repl}}$ individuals are selected randomly from this temporary population.

4. The $N_{\text{repl}}$ selected individuals produce the same amount of offspring by means of cross-over with a probability $p_{\text{cross}}$, or are otherwise just copied to the next generation. During cross-over, parts of their bit strings are swapped starting at a random place. The new strings, which contain parts of both parents, are stored in new
individuals which replace the ‘least-fit’ individuals of the previous generation, thereby keeping \( N_{\text{pop}} \) constant. Furthermore, some mutation may take place (with probability \( p_{\text{mut}} \)) in which bits are set randomly to 0 or 1. The fitness of the new individual is computed in the same way as in step 2.

5. All individuals are again sorted according to fitness. The algorithm repeats the previous step until the desired number of generations has been reached, or no significant improvement is reached anymore.

6. Finally, the parameters for each individual are stored in a file, which can be used to recreate the models or as input for a new series of generations.

We used the implementation of this algorithm in the GAlib genetic algorithm package, written in C++ by Matthew Wall at the Massachusetts Institute of Technology.

7.5.1. Test fit

In order to test this fitting procedure we attempted to infer the wind structure from the dynamic spectra shown in Fig. 7.8. The parameter space was limited to a maximum of 5 CIRs with \( \sigma_{\phi} \in [0.01, 0.2] \), \( \sigma_{r} \in [0.01, 0.5] \) and \( \Delta r_{\text{max}} \in [0.2, 5] \), \( \phi_{1} \in [0, 1.2] \), \( r_{0} \in [1, 1.4] \). For GA parameters we chose \( N_{\text{pop}} = 500, N_{\text{rep}} = 450, p_{\text{mut}} = 0.68, n_{\text{bit}} = 6 \).

For the fitness computation (see Eq. 7.17) we used \( \zeta = 0.5 \) and the fit was performed for the \( S = 0.2 \) model. The resulting dynamic spectrum is shown in Fig. 7.13, which has a fitness \( f = 7.9 \times 10^{-3} \) assuming \( \sigma_{ij} = 1 \) (see Eq. 7.16). Fig. 7.14 shows the fitted structure.

After the fit clearly converged too early (after 100 generations) with \( \langle D \rangle > 9.9 \times 10^{-5} \) and \( \sigma_{D} = 7.3 \times 10^{-5} \). This means that the average difference between the fitted and real flux is 0.01. Considering the lack of noise and absorption amplitudes of only about 0.05 we would expect \( \langle D \rangle \approx 10^{-6} \) for a good fit. Although the model dynamic spectrum qualitatively follows the “observed” dynamic spectrum there is still much to improve. Narrowing the parameter space by allowing only 3 CIRs or even increasing the population size to 5000 (with \( N_{\text{rep}} = 4500 \)) did not improve anything. The value of \( f \) always converges to about 8000 regardless of the dimension or population size. An explanation for this could be a too weak dependence of the solution on the parameters. Solutions become especially insensitive to \( \phi_{1} \) when \( \phi_{1} \gtrsim 0.5 \) and to most parameters when \( r_{0} \lesssim 1 + \sigma_{r} \). Therefore the parameter space can contain many plateaus with very broad extrema in which the algorithm can easily get trapped. We expect that defining a more meaningful parameter space is essential for getting good solutions.

Application of this fitting procedure to the observations of \( \xi \) Per did not result in a satisfactory solution.

7.6. Discussion and future work

The modeling of variations in O-star Hα profiles provides more insight in the possible location and shape of corotating wind structures. Due to the rotation of the star, each time a different part of the wind is probed, so that a picture of the wind structure emerges. A complication is that different regions in the wind simultaneously contribute to the Hα line formation and overlap in (projected) velocity. This contrary to the ultraviolet resonance lines where mainly changes in the blue-shifted P-Cygni absorption contribute to the observed variability, which can be uniquely identified with a location (in the absorbing column in front of the star) and a velocity (projection effects can be neglected).
One of the main conclusions is that most structures produce features which move from blue to red because of the relatively large tangential velocity due to angular momentum conservation. Absorption which moves outwards to more negative velocities is only produced when the perturbation starts further out in the wind, such that $r_0 > 1 + \sigma_t$. The variability due to changes in wind emission is even more dominated by stellar rotation and projection effects, because also the regions outside the line of sight contribute to the line formation. Furthermore, when only a blue-ward moving feature is observed in the dynamic spectrum, the structure must have a strong curvature $\phi_1 \gtrsim 0.7$. This seems to be the case for the O7.5 III star $\xi$ Per. For the supergiants we cannot say much about the curvature.

The curvature we derive is much larger than the value $\phi_1 \simeq 0.025$ used in the hydrodynamical model by Cranmer & Owocki (1996) and our analysis of the UV DACs in $\xi$ Per (De Jong et al. 2000a, Chapter 2). With such a large value of $\phi_1$ the CIR structure would spiral around the star many times, which does not seem to be consistent with the time evolution of the DACs. We must add, though, that often multiple (up to 4) DACs are detected in one spectrum, indicating that, in the interpretation of the CIR model, a CIR might cross the line of sight more than once. Anyway, given the fact that H$\alpha$ must be formed close to the star, we would have expected to see in H$\alpha$ the variations induced by only the “head” of the CIRs (which would resemble the “bar” like structures shown in Figs. 7.6(a) to 7.6(c)), and not by a large part of the “tail” of the CIR.

One of our best guesses for the wind structure in $\xi$ Per based on the variations in the H$\alpha$ profile is shown in Fig. 7.9. As said, the curvature of the CIRs is larger than expected on the basis of the interpretation of the DACs in UV resonance lines, but a spiral-shaped wind structure starting at some distance above the stellar surface is needed to explain the blue-shifted absorption events in H$\alpha$. We note that these absorption events are directly linked with the observed variations in the subordinate N IV line at 1718 Å and the DACs in the Si IV resonance lines at 1400 Å (De Jong et al. 2000a, Chapter 2) which are formed further out in the wind. We obtained better agreement with the observed time series by adding two blobs very close to the surface of the star (see Fig. 7.9).

One could speculate that the latter blobs represent the increase in mass flux above a bright spot on the stellar surface, one of the main ingredients of the CIR model. The wind structure displayed in Fig. 7.9 would be consistent with two locations on the stellar surface with an increased mass flux and the two resulting CIRs that are formed further out in the wind where the slow wind stream is overtaken by the faster wind coming from below. In our view, the two regions with a higher mass loss rate would correspond to the poles of an oblique, dipolar magnetic field connected to the star.

Altogether we conclude that it is most likely that the H$\alpha$ variations are caused by the same CIRs as the UV DACs (thus with the same curvature), but our qualitative analysis of the H$\alpha$ observations is not conclusive about this. Given the many possible solutions a quantitative fitting procedure is needed which also searches the parameter space simultaneously in many different directions. The Genetic Algorithm which we tested for this purpose only marginally fitted our test structure and we were not yet able to fit the real observations of $\xi$ Per. Improvements are needed in order to prevent an early convergence to a local minimum. Also, in stead of an empirical fit of the “undisturbed” H$\alpha$ profile, parameters like $\beta$, $\kappa$ and $S$ should be calculated self-consistently. Obviously, one should take into account the
continuity equation as well, which is not trivial. Finally, an extension to three dimensions is required in order to properly account for the (partial) coverage of the star by structures in the wind.

Acknowledgements. We thank Huiib Henrichs and Stan Owocki for their comments and help. JDJ acknowledges support from the Netherlands Foundation for Research in Astronomy (NFRA) with financial aid from the Netherlands Organization for Scientific Research (NWO) under project 781-71-053.

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Chapter 8

Nederlandse samenvatting

Dit proefschrift gaat over onderzoek naar de oorzaak van de variabiliteit in de wind van zware, hete sterren. Deze oorzaak is ondanks 15 jaar intensief speurwerk nog steeds niet gevonden, alhoewel er thans duidelijke ideeën zijn waar gezocht moet worden.

Onderzoek naar dit merkwaardige gedrag van deze sterren is belangrijk vanwege verschillende redenen. Deze sterren zijn de helderste sterren in het heelal. Ze worden ge-classificeerd als spectraaltype O en B. Vergeleken met de Zon is de massa van deze sterren 8 tot 100 maal groter, en bijgevolg is de druk en daardoor ook de temperatuur in het centrum van zo'n ster veel hoger. Onder deze omstandigheden verlopen de kernfusiereacties op een andere manier (de waterstofverbranding gaat via de CNO cyclus) en tevens ook veel sneller, zodat deze sterren onderruim en zijn tot een miljoen maal meer licht uitstralen dan de Zon. Dit heeft tot gevolg dat ze ook veel sneller door hun brandstofvoorraad zijn en slechts tot enkele miljoenen jaar kunnen bestaan, in tegenstelling tot de Zon met een levensduur van 10 miljard jaar. Er komen relatief slechts weinig van deze sterren in ons melkwegstelsel, maar dankzij hun grote lichtkracht zijn ze tot vele honderden lichtjaren met het blote oog te zien. Zo behoren bijna alle heldere sterren in het opvallende sterrenbeeld Orion tot de O en B spectraaltypen. Dit soort sterren markeren de spiraalarmen die in veel melkwegstelsels zichtbaar zijn. Ze worden namelijk geboren in deze spiraalarmen (die een relatief hogere dichtheid hebben), en hebben zich vanwege gedurende hun relatief korte levensduur nauwelijks van hun geboorteplaats kunnen verwijderen. De intense straling van deze sterren verhit het omringende gas en doet het oplichten. De eigenschappen van het interstellair medium wordt dan ook in hoge mate hierdoor bepaald.

Eén van de opvallendste fenomenen bij deze sterren is de sterke sterrenwind, waarvan nu duidelijk is dat die gedreven wordt door de druk van de uitgezonden straling. Op hun weg naar buiten botsen de in de kern geproduceerde lichtdeeltjes (fotonen) talrijke malen met de materie in de mantel en in de atmosfeer van de ster. In de atmosfeer dragen ze daarbij zoveel impuls aan de gasdeeltjes over, dat deze worden weggeblazen. Hierdoor wordt de zogenaamde sterrenwind in stand gehouden. Ook de Zon heeft zo'n wind (de zonnwind), waarvan is gemeten dat de snelheid in de buurt van de Aarde 500 km/s bedraagt. Het magneetveld van de Aarde beschermt ons tegen deze continue stroom van geladen deeltjes. In totaal verliest de Zon meer dan 600 miljoen kilogram materiaal per seconde. Dit lijkt misschien veel, maar de O en B sterren stoten per seconde bijna een miljard maal zoveel materie uit, waarbij uiteindelijk snelheden tot 3000 km/s bereikt kunnen worden.

Door dit intense massaverlies kunnen ze gedurende hun levensduur van doorgaans enkele miljoenen jaren wel de helft van hun massa kwijtraken. Uiteindelijk leidt dit ertoe dat de gehele mantel van de ster wordt weggeblazen en alleen de hete sterkern nog overblijft als een Wolf-Rayet ster is geworden. Aan het einde van zijn leven oploft de ster in een supernova explosie. Hierbij komen alle zware elementen die bij de kernfusie zijn gecreëerd in de interstellaire ruimte terecht en kan de schokgolf het ontstaan van nieuwe sterren tot gevolg hebben. De kern imploedeert daarbij tot een neutronenster of een zwart gat. Het is duidelijk dat de sterrenwind een grote invloed heeft op de evolutie en de interstellaire omgeving van deze zware sterren. Het is daarom belangrijk om dit massaverlies goed te bestuderen.

Met behulp van spectroscopische technieken is het mogelijk de atmosfeer en wind van een ster te onderzoeken. In de spectra zijn de spectraallijnen te identificeren van de elementen die op en rond de ster voorkomen. Deze lijnen hebben bovendien een karakteristieke vorm, waaruit af te leiden is hoe het materiaal rond de ster is verdeeld en welke snelheden het ten opzichte van de waarnemer beweegt. Door het Doppler-effect wordt de golflengte van de straling van materiaal dat op ons afkomt naar het blauw (korte golflengte) en materiaal dat van ons af beweegt naar het rood (langere golflengte) verschoven. Dit is vergelijkbaar met de verandering in toonhoogte van een passerende auto met sirene. Bij nadering is de toon van de sirene hoger (d.w.z. "blauwer") dan bij verwijdering. Uitstromende sterrenwind is op deze manier goed herkenbaar in spectraallijnen. Het materiaal dat zich namelijk voor de ster zelf bevindt beweegt naar ons toe, en zal het licht van de ster absorberen, met als gevolg dat er een blauw verschoven absorptielijn te zien is. De overige delen van de
uitstroomende wind zal een symmetrische emissielijn geven. Superpositie van beide lijnen geven zo een karakteristieke vorm, P Cygni lijnen genoemd, naar de eerste ster waarin nagenoeg alle spectrallijnen deze vorm hebben. Alleen lijnen die in de sterrenwind voorkomen hebben deze vorm. De lijnen die in de atmosfeer gevormd worden zijn niet verschoven en zijn symmetrisch. Door een aantal geschilde lijnen gelijktijdig te bestuderen kunnen we aldus de hele sterrenwind in kaart brengen. Het blijkt dat bij O en B sterren de sterrenwindlijnen in het ultraviolette (UV) deel van het spectrum voorkomen, dat niet vanaf de grond waarneembaar is. Voor de bestudering hiervan is dus een spectrograaf buiten de dampkring nodig.

Sinds de UV spectra van O en B sterren met satellieten zijn waargenomen, is bekend dat de P Cygni profielen van deze lijnen onverwacht sterke variaties vertonen. Dankzij de vele tijdsrekenen die zijn verkregen met de International Ultraviolet Explorer (IUE) satelliet (operationeel van 1978 t/m 1996), is duidelijk geworden dat deze variaties een zekere regelmaat vertonen. Het proefschrift van L. Kaper (1993) is hier geheel aan gewijd. Het gaat om extra absorptie, gesupereerd op de P Cygni lijnenprofielen, die bij een lage snelheid beginn en daarna gedurende een dag tot enkele dagen naar hogere snelheden optuigt. Soms zijn er meer- dere van dergelijke absorptiecomponenten aanwezig. Verder is na veel studie, o.a. vanuit Amsterdam, gebleken dat bij sterren met een hoge rotatiesnelheid dit patroon zich naar een kortere tijd min of meer herhaalt. Bij langzamaan draaiende sterren speelt zich de verschuiving veel langzamer af. Zie fig. 1.1 voor een paar voorbeelden.

Uit veel theoretische studies blijkt dat de uitleg voor dit fenomeen moet worden gezocht in het bestaan van spiraalvormige patronen van verhoogde dichtheid in de wind die niet aan de ster meeroteren (zogenaamde "Corotating Interaction Regions" ofwel CIRs). Men denkt dat deze patronen ontstaan doordat op sommige plaatsen boven het steroppervlak de wind iets minder (of juist meer) wordt versneld dan elders. De langzame stromen worden dan verderop in de wind ingehaald door de snellere wind. Hierdoor ontstaat een spiraalvormig patroon die de waarnemer eens per rotatie voor de ster langs ziet bewegen (zie fig. 1.2). Dit model verklaart de belangrijkste eigenschappen van de variaties, met name dat bij snelleroterende sterren de tijdschaal waarop de kleine absorptie patronen verschuiven inderdaad korter is. Blijkbaar zit er dus een of andere structuur op het oppervlak van de ster, die dit alles veroorzaakt, en waar de oorsprong van de veranderlijkheid van de sterrenwind gezocht moet worden.

Zoals de titel aangeeft, concentreert het onderzoek in dit proefschrift zich op oorsprong van deze veranderingen. Bij de aanvang van dit werk was niet goed bekend tot hoe dicht bij de ster de spiraalpatronen nog te traceren zijn en wat er nu precies gebeurt aan het oppervlak van de ster. Daarom hebben we eerst tegelijkertijd vanuit de ruimte en vanaf de aarde de sterrenwind waargenomen in het UV en optische gedeelte van het spectrum. Om de omstandighe-
tijd bij het volgen van ξ Persei verhindert. Deze tijd hebben we gebruikt om het magneetveld op een andere ster te zoeken: de B ster β Cephei, waarvan we al 15 jaar vermoedden dat er een behoorlijk veld aanwezig moest zijn. Het was zeer verheugend dat we inderdaad zeer nauwkeurige magneetveldmetingen konden verkrijgen, waarmee dus het veld (van 100 gauss) is gemeten van de heetste ster tot nu toe (zie hoofdstuk 6). De zeer regelmatige veranderingen in de wind van deze ster, die vroeger gemeten waren, konden geëxtrapolleerd worden, en een redelijk goed model kon worden opgesteld, waarmee ook de röntgenstraling van deze ster verklaard kon worden. Hiermee is in ieder geval aangetoond dat hete sterren een magneetveld kunnen bezitten, en dat de sterrenwind daarvan grote invloed ondervindt.

Tot slot hebben we in hoofdstuk 7 op basis van de CIR theorie een model gemaakt voor de variaties in de optische spectrumlijnen. Dit model blijkt de waarnemingen vrij goed te kunnen beschrijven, alhoewel het slechts in weinige gevallen getoetst kon worden. Kortom, uit dit proefschrift is te concluderen dat de CIR theorie redelijk goed werkt en dat de verstering aan de basis van de wind waarschijnlijk door een magneetveld wordt veroorzaakt. Om dit te bevestigen zullen betere metingen met grotere telescopen en meer gevoelige instrumenten nodig zijn.
Dankwoord

Graag maak ik van deze gelegenheid gebruik om iedereen te bedanken die direct of indirect heeft bijgedragen aan het tot stand komen van mijn proefschrift.

In de eerste plaats bedank ik Huib Henrichs bijzonder voor zijn uitstekende begeleiding en prima samenwerking tijdens mijn onderzoek. Ik heb van hem alle vrijheid gekregen die ik wenste om naar eigen inzicht dit onderzoek uit te voeren, maar zijn goede adviezen waren voor mij heel belangrijk om de goede richting en het overzicht te blijven houden. Ik heb al onze gesprekken, zowel binnen als buiten de sterrenkunde, altijd zeer gewaardeerd. Ook heeft hij veel bijgedragen aan mijn vaardigheden in het geven van lezingen, schrijven van artikelen en netwerken. Ook bedank ik Huib en Bep nog voor hun gastvrijheid. Ik kijk nog met veel plezier terug naar alle etentjes en reizen.

Ook Lex Kaper wil ik hartelijk bedanken voor zijn hulp, vooral tijdens de laatste twee jaar. Hij heeft mij veel gestimuleerd om projecten op tijd af te ronden en alles goed op te schrijven. Vooral in de laatste maanden heeft ook hij heel wat avonden en weekenden besteed aan het nauwkeurig lezen van mijn manuscript (evenals Huib). Verder denk ik nog met veel plezier aan al onze wetenschappelijke discussies, zowel in Amsterdam als in München.

Ik bedank mijn promotor Ed van den Heuvel voor zijn goede adviezen tijdens onze werkbesprekingen. Ik wil ook alle overige commissieleden bedanken voor het lezen van mijn manuscript en constructieve commentaar. Vooral Rens Waters heeft dit heel nauwkeurig en uitgebreid gedaan.

Verder is natuurlijk ook de onderzoekservaring die ik tijdens mijn stage heb opgedaan, van belang geweest. Ik heb toen veel van Thomas Augusteijn en Jan van Paradijs geleerd over alle aspecten van het sterrenkundig onderzoek, van het doen van waarnemingen tot aan de interpretatie en het modelleren. Helaas kan Jan het eindresultaat niet meer lezen.

Tijdens mijn waarnemreizen naar ESO la Silla, Haute-Provence, Pic du Midi en La Palma heb ik ook met veel collega's prettig samengewerkt. Met name tijdens en na de MUSICOS commission run op La Palma heb ik goede ervaring opgedaan in instrumentatie en datareductie. Ik heb met Joana Oliveira, Rudolf le Poole, Bernard Foin en Eric Stempels prima samengewerkt en inspirerende discussies gehad. Ook tijdens de bewolkte nachten wisten wij ons goed te vermaken met o.a. tafeltennis en snooker.

Ik heb mij altijd prima thuis gevoeld bij de prettige en ongedwongen sfeer op het sterrenkundig instituut. Zo staan mij veel leuke herinneringen bij van alle praatjes en avondjes uit met mijn kamergenoot Rudy Wijnands. Ook Coen Schriijers was altijd welkom om een discussie eens flink uit te breiden met van alles en nog wat. Verder hebben Martin Heemskerk en Frank van der Hooft mij vaak gesteund met hun humor, scherpe vragen en opmerkingen die mij steeds weer mijn problemen en ideeën deden heroverwegen. Ik kijk natuurlijk ook met plezier terug op alle andere praatjes, instituut activiteiten waaronder schaatsen(?), borrels en etentjes. Ik kan in deze context onmogelijk alle namen noemen, maar ik denk o.a. aan Jacco van Loon, Peer Zaal, Alex de Koter, Paul Groot, Sacha Hony, Ciska Kemper, Jane Ayal, Erica Veenhof, Jeroen Homan en Dirk Edel.

Ik bedank mijn broer Rob en mijn goede vrienden Sander Schippers, Sander Bosman, Bobby van den Hoek en Hester Volten voor hun steun, gezelligheid, leuke gesprekken en andere sociale activiteiten. Ook was Rob vaak bereid om een paar huishoudelijke taken van mij over te nemen, als ik het weer eens veel te druk had. Verder is mijn vrijwilligerswerk bij volkssterrenwacht Copernicus in Haarlem altijd zeer aangenaam.

Tot slot bedank ik mijn ouders hartelijk voor alles wat zij voor mij hebben gedaan. Mede dankzij de telescoop, computers, boeken en ruimte om dingen uit te zoeken die zij mij hebben gegeven, raakte ik voorgoed in de ban van sterrenkunde en computers.