Some aspects of evolution and mass transfer in X-ray binaries

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ON THE ROTATIONAL HISTORY OF THE PULSARS IN MASSIVE X-RAY BINARIES

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ABSTRACT

We investigate whether the rotation of the presupernova component (i.e., helium star) in a massive close binary can be synchronized with its orbital motion during its lifetime (i.e., before the SN explosion). It is shown that for binary periods in excess of one day, the two suggested synchronization mechanisms for stars with radiative envelopes (radiative damping of the dynamical tide and tidally induced shear turbulence) yield time scales which are more than 10^4 times too long. Hence, the pulsars in massive close binaries are expected to be born in rapid rotation, like single pulsars. The spin-down of the slow pulsars cannot be explained by the propeller mechanism as the weak winds of their companion OB main-sequence stars result in spin-down time scales far larger than the lifetimes of the latter stars. However, the mechanism recently proposed by Kundt yields a sufficiently short spin-down time scale ($\sim 3 \times 10^5$ yr). When the companion star leaves the main sequence and increases its wind strength, the pulsar is expected to show a rapid spin-up, on time scales of order 10^2 to 10^4 yr, in good agreement with the observed spin-up rates of 3U 0900−40, 3U 1223−62, and GX 1+4.

Subject headings: pulsars — stars: binaries — stars: rotation — X-rays: binaries

I. INTRODUCTION

In a recent paper Lea (1976) concluded that the slowly pulsing X-ray sources were most probably born as slowly rotating neutron stars. She suggested that tidal forces in a presupernova (pre-SN) binary system would be sufficiently strong to lead to synchronization of the stellar rotation with the orbital period within the lifetime of the system. The implosion of the core of a synchronized slowly rotating star would then have produced a slow pulsar.

Here we calculate the tidal torque in the pre-SN binary. The torque appears to be at least 10^4 times too small to slow down the pre-SN component of the binary within its lifetime. Other slowdown mechanisms for the rotation of neutron stars in binaries are considered.

II. TYPES OF TIDAL FORCES POSSIBLE IN THE PRESUPERNOVA BINARY

We consider only massive close binaries, as three of the slow pulsars are associated with massive X-ray binaries [viz., 3U 0900−40 ($P = 283$ s), X Persei ($P = 13.94$ minutes), and 3U 1223−62 ($P = 11.6$ minutes, associated with WRA 977, cf. Hammerschlag-Hensberge, Zuiderwijk, and van den Heuvel 1976; Mauder 1976); two more slow pulsars are probably associated with B0 emission stars (see references in Maraschi, Treves, and van den Heuvel 1976)]. The pre-SN binaries most probably consisted of a helium star (the pre-SN star) and a massive main-sequence star (cf. Sutantyo 1974; van den Heuvel 1974). Such systems are the result of the first stage of mass exchange in a massive close binary, and can most probably be identified with Wolf-Rayet (WR) binaries, the helium star being the WR star (Paczyński 1971; Smith 1973; van den Heuvel 1973). The main point here is to examine whether the helium star in such a system can be brought into synchronous rotation before it explodes as a SN. At its birth the helium star is expected to rotate more rapidly than synchronous, because it is the contracted core of the initial primary star, left after the loss of the hydrogen-rich envelope to the companion. We consider here a binary consisting of an 8 $M_\odot$ helium star with an 18 $M_\odot$ main-sequence companion. These values closely resemble the masses of the components of a typical Wolf-Rayet binary such as V444 Cygni and HD 211853 (cf. Kuhi 1973) and are expected to be representative for the masses of the components of typical progenitor systems of a massive X-ray binary such as Cen X-3 or 3U 0900−40.

Helium stars have convective cores and radiative envelopes. In stars with radiative envelopes, tides are much less effective than in stars with convective envelopes. In the latter ones, turbulent viscosity is an effective mechanism to produce large tidal friction and provides synchronism in a short time. However, also for stars with radiative envelopes two rather effective mechanisms for the production of tidal friction have recently been proposed, namely the following.

a) Radiative Damping

Zahn (1975, 1976) has shown that in stars with radiative envelopes and convective cores, radiative damping acting on the dynamical tide can provide an efficient mechanism for producing tidal friction. An essential point is that the star should have a convective

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core. Zahn (1976) showed that this mechanism can explain the small orbital eccentricities observed in the short-period massive X-ray binaries. Most of the tidal friction is produced by this mechanism when the star is still near the zero-age main sequence, where its convective core is largest. The time scale for tidal synchronization by radiative damping is (cf. Zahn 1976):

$$t_s = \frac{100}{E_2}\left(\frac{R_1^3}{M_1}\right)^{1/2} \frac{I_1}{M_1 R_1^2} \frac{1}{q^2(1+q)^{5/6}} \left(\frac{a}{R_1}\right)^{8.5} \mathrm{s}, \quad (1)$$

where R_1 and M_1 are the radius and mass (in solar units) of the star that is to be synchronized (which we call the primary), a is the semimajor axis of the orbit, $q = M_2/M_1$ is the mass ratio of secondary and primary, I_1 is the moment of inertia of the primary star in solar units, and E_2 is its second tidal coefficient. E_2 is a stellar structure constant, analogous to the apsidal motion constant k_2; the coefficients E_n appear, in first approximation, to scale as $(R_c/R_1)^{2n+4}$, where R_c is the radius of the convective core of the star (Zahn 1976).

b) Shear Turbulence

A second mechanism, proposed by Press, Wiita, and Smarr (1975), is the generation of shear turbulence in the outer layers of the star by the traveling of the tidal bulge over the stellar surface. Although, on the basis of a linear stability analysis, Seguin (1976) has questioned the occurrence of this kind of turbulence, we assume for the sake of argument that this mechanism can operate efficiently. The equation for the synchronization time scale is in this case (Press, Wiita, and Smarr 1975)

$$t_s = \frac{75}{224} \frac{R_T \alpha}{K_\mu |\Omega - \omega|} (1 - e^2)^{9/2} \left(\frac{a}{R_1}\right)^9 \left(\frac{M_1}{M_2}\right)^2, \quad (2)$$

where R_T is the effective Reynolds number (assumed to be 20), K_μ is a structural constant defined by Press *et al.*, α is the square of the radius of gyration ($\alpha = I_1/M_1 R_1^2$), and Ω and ω are the orbital and rotational angular velocities, respectively. K_μ and α have values of about 0.001 and 0.1, respectively, for an 8 $M_\odot$ helium star.

III. APPLICATION TO HELIUM STAR BINARIES

We consider a system consisting of a 8 $M_\odot$ helium star and an 18 $M_\odot$ normal main-sequence star.

a) Radiative Damping

We used the evolutionary sequence of an 8 $M_\odot$ helium star computed by Savonije and Takens (1976) (the density structure of these models does not differ much from that of the models computed by Paczyński 1971). The stellar radius and core radius of a 8 $M_\odot$ helium star are approximately constant during core helium burning and are 0.75 $R_\odot$ and 0.3 $R_\odot$, respectively. It follows that during core helium burning E_2 is smaller than 10^{-5}. This yields a lower limit to the synchronization time scale of

$$t_{s_{\min}} = 8 \times 10^6 a^{17/2} \approx 6 \times 10^8 P^{17/3}\ \mathrm{yr}. \quad (3)$$

(a in $R_\odot$, P in days). Equation (3) shows that, in order to reach synchronous rotation within the lifetime of the helium star (7×10^5 yr; because of neutrino losses the lifetime after core helium burning is negligible), P should be less than 0.3 days. Or, if one adopts $P = 3$ to 5 days (which seems a characteristic value for short-period WR binaries as well as for progenitors of X-ray binaries like 3U 0900−40), one finds that the synchronization time scale is more than 5×10^5 times the lifetime of the helium star. Even for the WR binary with the shortest known period (CQ Cephei, $P = 1.64$ days) the synchronization time is still 10^4 times longer than the stellar lifetime.

b) Shear Turbulence

Application of equation (2) to the same system, with the above mentioned values of K_μ, α, and R_T, and $|\Omega - \omega|/\Omega \approx 20$ (i.e., the angular velocity of rotation 20 times faster than the orbital motion, as follows from the ratio of the moments of inertia of the stellar core at the end of hydrogen burning, and the helium star) yields a lower limit

$$t_{s_{\min}} = 0.4 \times 10^{-3} a^{10.5}\ \mathrm{yr} = 1.2 \times 10^8 P^7\ \mathrm{yr}. \quad (4)$$

In order to achieve synchronism within the 7×10^5 yr lifetime of the helium star, P should here be shorter than 0.5 days. With $P = 3$ days, t_s becomes 4×10^5 times longer than the lifetime of the helium star. Hence, the occurrence of fully developed shear turbulence would not alter the picture: the helium star cannot be synchronized during its lifetime.

IV. OTHER SLOW-DOWN MECHANISMS FOR NEUTRON STARS IN MASSIVE CLOSE BINARIES

a) The Time Scale for Spin-down Inferred from Observations

From the above we conclude that the rotational evolution of the helium star will proceed completely independent of the presence of the companion star. Consequently, just as for a single star, the core collapse of an evolved helium star is expected to produce a rapidly spinning neutron star. Some of the slow pulsars are associated with B0 emission-line stars (notably X Persei and the two pulsing transient sources; cf. Stier and Liller 1976; Maraschi, Treves, and van den Heuvel 1976) which are in or close to the main sequence.

The mass of a main-sequence B0 star is around 20 $M_\odot$ (cf. Underhill 1966) and the main-sequence lifetime of such a star after the SN stage of its companion is expected to be less than some 5×10^6 yrs (cf. De Loore and De Greve 1975; De Loore, De Greve, and De Cuyper 1975). Assuming the slow pulsars to have been born with a spin period of less than 0.1 s, and the spin-down to have occurred exponentially, one finds that in order to have reached a period of 5 minutes in less than 5×10^6 yr, t_{sd} must have been shorter than 6×10^5 yr.

b) *The Propeller Mechanism*

The propeller mechanism (Illarionov and Sunyaev 1975) is not able to produce a spin-down on such a short time scale. Even if a very strong wind ($\dot{M} = 10^{-7} M_\odot \text{ yr}^{-1}$, $V \approx 500$-1000 km s^{-1}) is adopted for the companion star during its entire main-sequence life, the spin-down time scale with this mechanism is still of the order of 10^7–10^8 years (cf. Wickramasinghe and Whelan 1975; Fabian 1975).

This is the more serious, since it is well known from recent observations that normal as well as emission-line main-sequence stars of about spectral type B0 have only rather weak winds with mass loss rates not exceeding $\dot{M} = 10^{-8.5} M_\odot \text{ yr}^{-1}$ (Rogerson and Lamers 1975; Snow and Marlborough 1975). For such winds the propeller mechanism yields spin-down time scales of order 5×10^8 yr.

c) *Kundt's Mechanism*

Kundt (1976) has argued that for a neutron star spinning in a stellar wind, very large friction on the magnetosphere is to be expected as soon as the neutron star has sufficiently slowed down to allow the infalling wind matter to penetrate inside the velocity-of-light cylinder (this occurs for spin periods around 0.4 s, in winds with the above quoted strength for a B0 main-sequence star). A very large torque on the magnetosphere is then produced by the bending of the field lines in the transition region between the pulsar magnetosphere and the infalling stellar wind plasma. Because of this torque the infalling particles are accelerated to highly relativistic energies, and the pulsar will be slowed down very rapidly (cf. Davidson and Ostriker 1973).

The braking time scale with this mechanism is (Kundt 1976)

$$t_{sd} = 10^8 \Omega_0 I_{44.5} T_{34}^{-1} \text{ yr}\,, \tag{5}$$

where T_{34} and $I_{44.5}$ are the slowdown torque and the moment of inertia of the neutron star in units of 10^{34} and $10^{44.5}$ cgs, respectively. T_{34} is given by

$$T_{34} = 5.6 B_{12}^{2/7} a_{12}^{-12/7} \beta^{-24/7} \times [\dot{M}_p / 10^{-6}[M_\odot \text{ yr}^{-1}]]^{6/7}\,, \left(\frac{M_n}{M_\odot}\right)^{16/7} \tag{6}$$

where B and a are the surface magnetic field strength of the neutron star and the orbital radius, in units of 10^{12} gauss and 10^{12} cm, respectively; $\dot{M}_p$ is the mass loss rate from the primary star, β is the stellar wind velocity near the orbit, in units of 600 km s^{-1}, and M_n is the mass of the neutron star.

Adopting the observed stellar wind parameters for the B0 V main-sequence star τ Sco, i.e., $\dot{M} = 10^{-8.5} M_\odot \text{ yr}^{-1}$, and $v_w \approx 1000 \text{ km s}^{-1}$ (Rogerson and Lamers 1975), and $\Omega_0 = 2\pi$, $B_{12} = a_{12} = 1$, $M_n = M_\odot$, one obtains $t_{sd} = 3 \times 10^5$ yr. Consequently, neutron star companions to early-type main-sequence stars will very rapidly spin down to become very slow pulsars.

d) *The Equilibrium Spin Period and Anticipated Spin-up*

The spin down terminates when the pulsar has reached an equilibrium spin rate, at which for matter near the magnetospheric boundary the corotation velocity equals the Keplerian velocity around the neutron star (Davidson and Ostriker 1973). This spin period is for accretion from a wind given by (cf. Wickramasinghe and Whelan 1975; Illarionov and Sunyaev 1975)

$$\begin{aligned} P_{eq} &= 15 (B_{12})^{6/7} \left(\frac{1.5 \times 10^{-7} M_\odot \text{ yr}^{-1}}{\dot{M}_p}\right)^{3/7} \left(\frac{M_\odot}{M_n}\right)^{11/7} \\ &\quad \times \left(\frac{a}{20 R_\odot}\right)^{6/7} \left(\frac{V_w}{10^3 \text{ km s}^{-1}}\right)^{12/7} \text{s} \\ &= 1.6 (B_{12})^{6/7} \left(\frac{M_\odot}{M_n}\right)^{5/7} \left(\frac{10^{17} \text{ g s}^{-1}}{\dot{M}_a}\right)^{3/7} \text{s}\,, \end{aligned} \tag{7}$$

where V_w is the wind velocity relative to the neutron star and $\dot{M}_a$ is the accretion rate onto this star (note that the constant of proportionality has been corrected). With the above quoted wind parameters for the B0 V star τ Sco and $B_{12} = 1$, $M_n = M_\odot$, $a = 50 R_\odot$ (like for 3U 0900−40), one obtains $P_{eq} = 172$ s. With $\dot{M} = 10^{-9} M_\odot \text{ yr}^{-1}$ one obtains 282 s, i.e., practically equal to the present spin rate of 3U 0900−40.

Consequently, the spin rates of the slow pulsars seem to agree excellently with the equilibrium spin rate expected for a neutron star in the weak wind of a normal early-type main-sequence star.

An interesting point noticed by Wickramasinghe and Whelan is that the accretion rate of about 10^{-10} to $10^{-9} M_\odot \text{ yr}^{-1}$ derived from the X-ray luminosity of 3U 0900−40 (P = 283 s) yields—with $B_{12} = 1$, $M_n = M_\odot$—in equation (7) a value $P_{eq} = 2.7$–7.3 s; i.e., apparently, this source was spun down in a wind much weaker than the present one of its companion.

On the other hand, for Cen X-3 and SMC X-1 one derives, from the accretion rates of $10^{-9} M_\odot \text{ yr}^{-1}$ and $10^{-7.5} M_\odot \text{ yr}^{-1}$, respectively (inferred from their X-ray luminosities), that these sources should have equilibrium spin periods of about 2.7 s and 0.7 s, respectively (again for $B_{12} = 1$ and $M_n = M_\odot$). Taking into account that B_{12} as well as M_n may be slightly different from the adopted values, one must conclude that these values agree well with the observed pulse periods of these sources (4.84 s and 0.71 s, respectively). It therefore seems very likely that Cen X-3 as well as SMC X-1 are presently rotating near equilibrium with the stellar winds of their companions.

The fact that both sources show extended low states fits well with such a situation, since for neutron stars spinning near equilibrium, the accretion of angular momentum may cause the accretion to be shut off temporarily from time to time. On the other hand, 3U 0900−40 is expected to be continuously spinning up, as its present spin period is much longer than the equilibrium period for the (strong) wind of its companion. The accretion wake angle of 3U 0900−40 as well as the Hα profile of its companion star HD 77581 indicate wind velocities of between 200 and 400 km s^{-1} near the orbit (cf. Charles *et al.* 1976; Zuiderwijk, van den Heuvel, and Hensberghe 1974; Bessell, Vidal, and Wickramasinghe 1975). Adopting an accretion rate of 10^{-9} $M_{\odot}$ yr^{-1}—comparable with the X-ray luminosity—one can, for each adopted wind velocity, calculate the wind density. Inserting the resulting values into Fabian's (1975) equation for the spin-up time scale, one finds that with $V_w = 400$ km s^{-1}, 3U 0900−40 should spin up from $P = 283$ s to $P = 60$ s in about 2×10^4 yr. With $V_w = 300$ km s^{-1} and 200 km s^{-1} the spin-up time is 7×10^3 and 4×10^2 yr, respectively.

Consequently, a rapid spin-up of 3U 0900−40 is to be expected. Indeed, a spin-up time scale of around 10^4 yr has recently been reported (Rappaport and Joss 1976).

The sources 3U 1223−62 and GX 1+4 show an even faster spin-up, both on time scales of order 10^2 yr (Swank *et al.* 1976; Becker *et al.* 1976), corresponding to wind velocities near the orbit of about 200 km s^{-1}. This is indeed about the velocity inferred from the Hα profile of the probable companion of 3U 1223−62, which is the B1.5 Ia star WRA 977 (Hammerschlag-Hensberghe, Zuiderwijk, and van den Heuvel 1976; Mauder 1976).

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