'ev. Mexicana Astron. Astrof., 21, 309-313 (1990)

ROTATIONAL EVOLUTION AND MAGNETIC FIELD OF Ap STARS

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RESUMO. Propõe-se que o campo magnético de estrelas Ap pode ser gerado pelo mecanismo de dínamo na base do envelope convectivo, e transportado para a superfície pela instabilidade de boiamento na fase de Hayashi. Campos magnéticos superficiais observados permitem estimar uma perda de momento angular durante a fase pré-Seqüência Principal compatível com as observações. Estrelas A normais, que têm rotação rápida, não mostram campos magnéticos superficiais importantes e isto pode acontecer se uma protoestrela evolue para Seqüência Principal sem passar pela fase de Hayashi.

ABSTRACT: It is proposed that the magnetic field of Ap stars may be generated by the dynamo mechanism at the base of the convective envelope, and transported to the surface by the instability of buoyancy in the Hayashi phase. Observed surface magnetic fields allow to estimate a loss of angular momentum during the pre-Main Sequence phase compatible with the observations. Rapidly rotating normal A stars do not show important surface magnetic fields and this may occur if a protostar evolves to Main Sequence skipping the Hayashi phase.

Key words: HYDROMAGNETICS — STARS-PECULIAR A

. INTRODUCTION

Observations point out that solar dipolar magnetic field is of ≈ 1 G. resumably this field is generated by dynamo mechanism. Chemically peculiar Ap ›tars are of spectral type for which the magnetic dynamo cannot be invoked in he Main Sequence. Generally they have surface magnetic fields of ≈ 10^3 G and a ›lower rotation than the normal A stars. Phenomena of Ap stars have been nterpreted through the model of oblique rotator (Khoklova, 1985). The majority ›f authors believe that magnetic field of Ap stars is fossil (Borra 1982, ›olginov 1985). A statistic property, though not definitely established, is a ossible anti-correlation between the strength of surface field and rotational ›elocity (Mestel, 1975).

The great differences between the Ap and A stars will be analysed ›ith regard to the surface magnetic field and angular momentum. It is assumed hat the angular momenta of A and Ap stars do not differ significantly nmediately after the collapse. Measurements of pseudo angular momentum (J = ›RV) of Main Sequence A (Kraft, 1970) and Ap stars (Khokhlova, 1985) show that stars of the latter type suffered considerable loss of angular momentum perhap in the pre-Main Sequence phase.

Molecular clouds which give rise to A and Ap stars collapse in th subcritic regime (Mestel, 1987) as far as the magnetic field stays frozen Therefore the great magnetic field difference between A and Ap stars must b explained by some process which acts in the pre-Main Sequence phase.

Larson (1969, 1972) analysed the evolution of molecular clouds int protostars with 2 - $3M_{\odot}$. The size of the convective envelope is larger c smaller, according to the condensation degree in the beginning of collapse ha been smaller or greater.

It will be assumed that the magnetic field of Ap stars is produced b dynamo mechanism at the base of the convective envelope during the Hayas phase. A simple dynamo model which will be used in this work is that one fro Durney and Robinson (1982) (DR model for brevity). The amplification time c the magnetic field is made equal to the rise time of the magnetic flux tubes u to the surface by magnetic buoyancy (Acheson,1978).

2. EQUATIONS

Equations are in c. g. s. units unless stated otherwise.

a) Magnetic Field at the Base of the Convective Envelope

After the DR model, the magnetic amplification time T_A through the α effect is

$$T_A = (L/\alpha\Delta\Omega)^{1/2}$$

where L is the scale height for pressure; $\alpha = C_1L^2\Omega/R_c$, $\Delta\Omega = C_2(L/R_c)^2\Omega$, R_c i the internal radius of the convective envelope (or the radius of radiativ core); Ω is the angular velocity; C_1,C_2 are empirically adjusted constants.

The rise time of the magnetic flux tube by buoyancy is

$$T_r = L/U$$

where

$$U = V_A{}^2(3Q/8)(X/L)^2/V,$$
$$Q = \ln(4/N_A)-0.077,$$
$$N_A = (9Q/8)(X/L)^3(V_A/V),$$
$$V = \ell/2(g/T)^{1/2}(\nabla\Delta T)^{1/2}$$
$$F_c = \ell^2/4C_p\rho(g/T)^{1/2}(\nabla\Delta T)^{3/2}$$

$X = L/2$ is the radius of the flux tube; V_A and V are the Alfvén and convectiv velocity respectively; N_A is magnetic Reynolds number; F_c is the convectiv energy flux; $\ell \approx R_\star - R_c$ is the mixing length; $R_\star$ is the radius of the star; is gravitational acceleration; C_p is the specific heat for constant pressur and ρ is the mass density; T is the temperature.

At the base of the convective envelope the magnetic field B_b i obtained by making $T_A = T_r$,

$$B_b = D(\rho VL^{5/2}/QR_c{}^{3/2})^{1/2}$$

For stellar mass of 2.5 $M_{\odot}$, D = 0.01913 for $R_\star$ = 5 $R_{\odot}$. In the above equation R is a free parameter. A differential rotation exists in a shell at the bottom o the convective envelope.

b) Instability of Buoyancy

The first articles on this instability have been written by Parke (1975, 1977a). Later Parker (1977b) included the influence of rotation an finally Acheson (1978) included the effects due to magnetic and thermal diffusivities and viscosity.

Here the perturbations will be considered non-axisymmetric; $V_A^2 \ll \Omega^2 R^2$; thermal and magnetic difusivities null and viscosity small. The instability criterion (Acheson, 1978) was applied for the convective regions of the Ap protostars. The stellar structure was defined by a polytropic model (γ=5/3) with the Larson's boundary condition (1969). For the radial direction R of a spherical coordinate system, the instability criterion is

$$-\left(\frac{g}{\gamma a^2} - \frac{2}{R}\right) V_A^2 \frac{\partial F}{\partial R} > \frac{m^2 V_A^2}{R^2}$$

Here a is the acoustic velocity; F=ln(B/ρR); the azimuthal wave number m=1. The criterion of instability for direction z is

$$-\frac{g}{\gamma a^2} V_A^2 \frac{\partial F}{\partial R} > \frac{m^2 V_A^2}{\cos^2\alpha R^2}$$

where α is the latitude angle. These instability criteria give a limiting gradient of magnetic field and allow to calculate a minimum field B_s at the surface, given the value of B_b determined by the DR model.

c) Loss of Angular Momentum

Hartmann (1986) elaborated for T Tauri stars a model of mass loss $\dot{M}$ excited by Alfvén waves:

$$\dot{M} \sim 10^{-14} M^{-1.5} R_*^{3.5} F_5^2 B_s^{-2} \quad M_\odot \ \ yr^{-1}$$

Here M and R_* are respectively the mass and radius of the star measured in solar units; the Alfvén wave flux F_5 at the stellar surface is in unit of 10^5 erg $cm^{-2}s^{-1}$; B_s is the surface magnetic field in Gauss. The corresponding rate of angular momentum loss $\dot{J}$ is given by

$$\dot{J} = \dot{M} R_A^2 \Omega .$$

where R_A is the Alfvén radius.

3. CALCULATIONS AND RESULTS

Calculations have been made adopting a stellar mass of 2.5 $M_\odot$ and a stellar radius of 5 $R_\odot$. For such a protostar the base of the convective envelope, R_c, cannot be deeper than ≈ 0.56 R_*. Physical meaning of this limit is the cessation of the buoyancy and infinite amplification of the field there. In all Figures presented below R_c is in the range 0.56 $R_* < R_c < R_*$.

(i) Using the DR model for a protostar with 2.5 $M_\odot$ and radius of 5 $R_\odot$, T_A and T_r can be computed as a function of R_c. The angular velocity at the top of the convective envelope was fixed at the limiting value defined by the critical rotational velocity of break-up. The results for B = 10^5, 10^4, 10^3 G are shown in Figure 1.

(ii) Figure 2 shows B_b as a function of R_c.

(iii) Figure 3 illustrates the minimum surface field B_s for five positions of the convective envelope base. As deeper is this base, more intense is B_s.

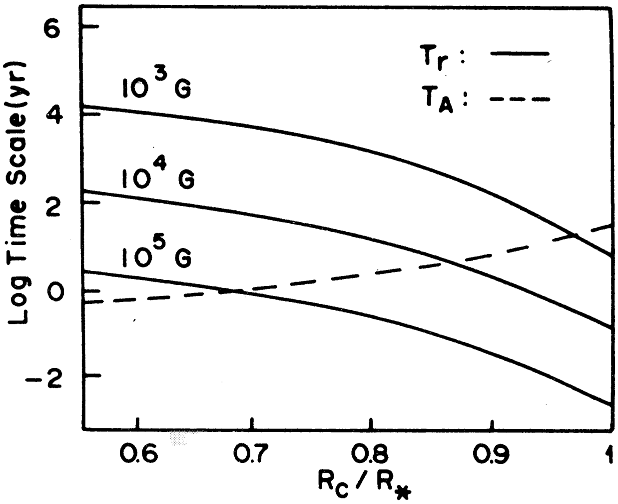

Figure 1. Magnetic field amplification time, T_A, and magnetic flux tube rise time, T_r, as function of the radial position of the base of convective envelope. Curves for T_r have been computed for a constant value of magnetic field at the base of the convective envelope, indicated in the Figure.

(vi) Using surface magnetic field ≈ 10^3 G, the mass loss can be estimated according to the Hartmann's model (1986). The corresponding loss of angular momentum is 7×10^{52} (g cm^2/s) during the pre-Main Sequence phase. Therefore the magnetic rotational braking time scale is 10^4 – 10^5 years.

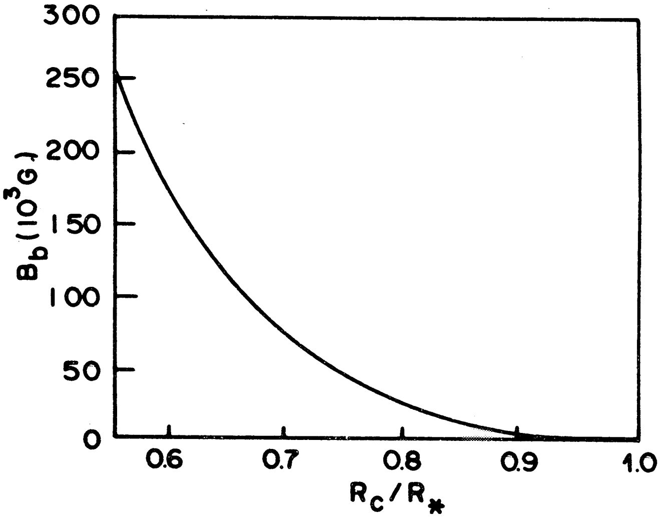

Figure 2. Amplified magnetic field, B_b, as a function of the radial position of the base of convective envelope.

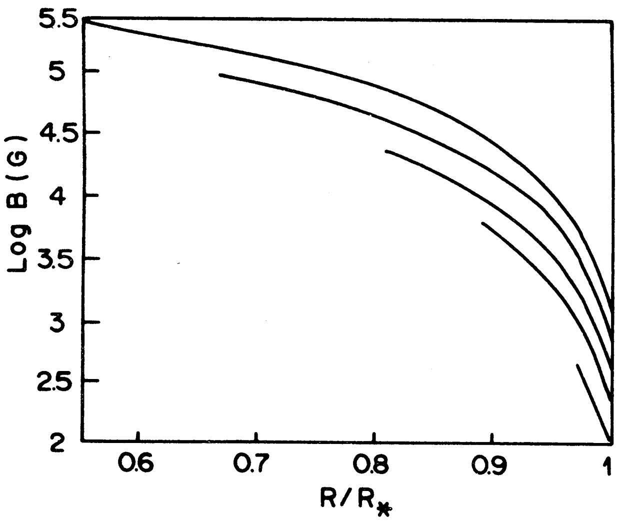

Figure 3. Magnetic field as a function of the radial distance, R, for 5 different depths of the convective envelope.

4. DISCUSSION

These preliminary results indicate that the magnetism of the Ap stars may be due to a dynamo process on the base of the convective envelope during the Hayashi phase. If such is the case the strength of surface field B_s decreases as the convective envelope becomes thinner. But the thickness of the convective envelope depends upon the initial conditions of the collapse. Then the variety of initial conditions may explain the existence of magnetic Ap stars and non magnetic A stars.

In the Main Sequence the instability of buoyancy is difficult to cur (Acheson,1978), so that the idea of buoyancy remains applicable only for e pre-Main Sequence phase.

The results obtained are in agreement with a claimed anti-correlation tween B_S and Ω.

Magnetic and rotational properties of stars of spectral types jacent to Ap stars can be considered. Certainly the O,B type stars do not mit the application of the present model because they do not pass through the yashi phase. Stars of spectral type F with color index (B-V) > 0.45 differ om the A and Ap stars because they show magnetic activity correlated with the tation (Wolff, 1987) and have convective envelope in the Main Sequence.

One of the authors, C.X., benefited from fellowships from CNPq roc. 840094-87.6) and CAPES during the preparation of this work.

FERENCES

heson, D.J. 1978, *Phil. Trans. Roy. Soc. London.*, 289A, 459.
rra, E.F. 1982, *Ann. Rev. Astron. Astrophys.*, 20, 191.
lginov, A.Z. 1985, in *IAU Colloquium No. 90, Upper Main Sequence Stars with Anomalous Abundances*, eds. C.R. Cowley, M.M. Dworetsky and C. Mégessier (Dordrecht: D. Reidel), p. 11.
rney, B.R., and Robinson, R.D. 1982, *Ap. J.*, 253, 290.
rtmann, L. 1986, *Fundamentals of Cosmic Physics*, 11, 279.
okhlova, V.L. 1985, *Sov. Sci. Rev. E Astrophys. Spase Phys.*, 4, 99.
aft, R.P. 1970, in *Spectroscopic Astrophysics.*, ed. G.H. Herbig (Berkeley: University of California Press), p. 385.
rson, R.B. 1969, *M.N.R.A.S.*, 145, 271.
rson, R.B. 1972, *M.N.R.A.S.*, 157, 121.
stel, L. 1975, in *IAU Colloquium No. 32, Physics of Ap Stars*, eds. W.W. Weiss, H. Jenkner and H.J. Wood (Vienna: Universitätssternwarte Wein), p. 1.
stel, L. 1985, in *Protostars and Planets II*, ed. D.C. Black (Tucson: The University of Arizona Press), p. 320.
rker, E.N. 1975, *Ap. J.*, 198, 205.
rker, E.N. 1977a, *Ann. Rev. Astron. Astrophys.*, 15, 45.
rker, E.N. 1977b, *Ap. J.*, 215, 307.
lff, S.C., and Heasley, J.N. 1987, *P.A.S.P.*, 99, 957.

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