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 dinamo 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-Sequência
Principal compatível com as observações. Estrelas A
normais, que têm rotação rápida, não mostram campos
magnéticos superficiais importantes e isto pode acontecer
se uma protostrela evolue para Sequência 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 \( \approx \ 1 \) G,
resumably this field is generated by dynamo mechanism. Chemically peculiar Ap
stars are of spectral type for which the magnetic dynamo cannot be invoked in
the Main Sequence. Generally they have surface magnetic fields of \( \approx \ 10^3 \) G and a
slower rotation than the normal A stars. Phenomena of Ap stars have been
interpreted through the model of oblique rotator (Khokhlova, 1985). The majority
of authors believe that magnetic field of Ap stars is fossil (Borra 1982,
Poligainov 1985). A statistic property, though not definitely established, is a
possible anti-correlation between the strength of surface field and rotational
ecllcity (Mestel, 1975).

The great differences between the Ap and A stars will be analysed
with regard to the surface magnetic field and angular momentum. It is assumed
that the angular momenta of A and Ap stars do not differ significantly
immediately after the collapse. Measurements of pseudo angular momentum \( (J \approx
1RV) \) of Main Sequence A (Kraft, 1970) and Ap stars (Khokhlova, 1985) show that
stars of the latter type suffered considerable loss of angular momentum perhaps in the pre-Main Sequence phase.

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

Larson (1969, 1972) analyzed the evolution of molecular clouds in protostars with $L \sim 3M_\odot$. The size of the convective envelope is larger or smaller, according to the condensation degree in the beginning of collapse hence been smaller or greater.

It will be assumed that the magnetic field of Ap stars is produced by dynamo mechanism at the base of the convective envelope during the Hayashi phase. A simple dynamo model which will be used in this work is that of Durisen and Robinson (1982) (DR model for brevity). The amplification time of the magnetic field is made equal to the rise time of the magnetic flux tubes 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

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

where $L$ is the scale height for pressure; $\alpha = C_1 L^2 \Omega / R_C$, $\Delta \Omega = C_2 (L/R_C)^2 \Omega$, $R_C$ is the internal radius of the convective envelope (or the radius of radiative core); $\Omega$ 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

$$U = V_A^2 (3Q / 8) (X / L)^2 / V,$$

$$Q = \ln(4 \pi / N_A) - 0.077,$$

$$N_A = (9Q / 8) (X / L)^3 (V_A / V),$$

$$V = L^2 (g / T)^{1/2} (\Omega T)^{1/2}$$

$$F_c = L^2 / 4C_p (g / T)^{1/2} (\Omega T)^{3/2}$$

$X = L / 2$ is the radius of the flux tube; $V_A$ and $V$ are the Alfvén and convective velocity respectively; $N_A$ is magnetic Reynolds number; $F_c$ is the convective energy flux; $I = R_\ast - R_C$ is the mixing length; $R_\ast$ is the radius of the star; $g$ is gravitational acceleration; $C_p$ is the specific heat for constant pressure; $\rho$ is the mass density; $T$ is the temperature.

At the base of the convective envelope the magnetic field $B_B$ is obtained by making $T_A = T_r$:

$$B_B = D (\rho V L^{5/2} / Q R_C^{3/2})^{1/2}$$

For stellar mass of $2.5 \ M_\odot$, $D = 0.01913$ for $R_\ast = 5 \ R_\odot$, in the above equation $R$ is a free parameter. A differential rotation exists in a shell at the bottom of 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;
$V_A^2 < \alpha^2 R^2$; thermal and magnetic diffusivities 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
($\gamma=5/3$) with the Larson's boundary condition (1969). For the radial direction $R$
of a spherical coordinate system, the instability criterion is

$$\frac{g}{\gamma a^2} \frac{aF}{R} - \frac{m^2 V_A^2}{R^2} > 0$$

Here $a$ is the acoustic velocity; $F = \ln(B/\rho R)$; the azimuthal wave number $m=1$. The
criterion of instability for direction $\theta$ is

$$\frac{g}{\gamma a^2} \frac{\partial F}{\partial \theta} - \frac{m^2 V_A^2}{\cos^2 \alpha \partial \theta^2} > 0$$

where $\alpha$ 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 $B_0$ determined by the DR model.

c) Loss of Angular Momentum

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

$$M = 10^{-14} M^{-1.5} R_x^{3.5} F_5^2 B_S^{-2} M_0 \text{ yr}^{-1}$$

Here $M$ and $R_x$ 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 $J$ is given by

$$J = MR_A^2 \Omega$$

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 $\approx 0.56 R_x$. 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 $0.56 R_x < R_C < R_x$.

(i) Using the DR model for a protostar with 2.5 $M_\odot$ and radius of 5
$R_\odot$, $T_A$ and $T_F$ 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 $B_0$ 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 $B_S$. 

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Figure 1. Magnetic field amplification time, $T_A$, and magnetic flux tube rise time, $T_r$, as a 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 $7 \times 10^{52}$ (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 between $B_S$ and $\Omega$.

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.

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