TURBULENT MIXING IN CIRCUMSTELLAR SPACE IN THE PRESENCE OF MAGNETIC FIELDS

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ABSTRACT. The magnetic field structure in the swept-up HII shell and the stellar wind region around an early type star is described. It is argued that the field does not become dynamically so important as to prevent self-similar flow. A condition under which reconnexion removes magnetic energy from the bubble is derived. It follows that the process of turbulent mixing between the HII region and the bubble works efficiently enough to cause severe departures from uniform expansion, even in the presence of magnetic fields.

1. Introduction

There have been many studies of classical wind-blown HII regions around OB stars and Planetary Nebulae Nuclei, where a thin, shocked shell consisting of an inner ionized and an outer neutral layer is held under pressure by a hot shocked stellar wind (HSSW) (e.g. Dyson and deVries, 1972; Castor et al. 1975; Breitschwerdt 1985). Although this model has been very successful in the past for our understanding of circumstellar dynamics, there are some severe discrepancies with observations that cannot be accounted for, such as the existence of rather extended HII shells (Dickel, 1974; Rodriguez et al., 1988), highly ionized species like OVI, NV (Jenkins and Meloy, 1974), soft X-rays (Burstein et al. 1977) and high velocity flows of the order of 100 km/s (López et al. 1989). Weaver et al. (1977) studied the formation of highly ionized species by global thermal conduction across the HSSW/HII interface. We adopt the view that any tangential global magnetic field, present either in the bubble or the ionized layer, suppresses conduction very efficiently. Instead we propose a model where mass and energy are exchanged between the two adjacent regions by the mechanism of turbulent mixing. A source of turbulence, which should exist in most ionization bounded, wind-blown HII regions, is the acoustic instability (Breitschwerdt and Kahn, 1988). It is generated by the amplification of sound waves travelling parallel to the HII/HSSW interface due to gas motions across the ionization front. We have shown that such waves can grow exponentially, thereby producing turbulence on a scale comparable to the thickness of the ionized layer.

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The process of turbulent mixing has been described in detail elsewhere (Kahn and Breitschwerdt, 1989); here we wish to concentrate on the structure of the magnetic field and its dynamical effects.

Observed wind-blown HII regions also show features which can be attributed to the structure of the various undisturbed circumstellar media, such as density gradients and clumpiness (for a recent review see Dyson, 1989). Our aim is, in a more fundamental sense, to reexamine the classical bubble picture and, from studying the flow properties, to point out that modifications have to be made.

2. Magnetic Field Structure in the Circumstellar Medium

In this section we summarize the structure of the magnetic field in the ambient medium and in the stellar wind region. For a more detailed discussion we refer to Kahn and Breitschwerdt (1989) and Breitschwerdt and Kahn (1989 (in preparation)). A typical example of a wind-blown HII region is NGC 6334(A) (Rodriguez et al. (1988)) and in what follows we shall use numerical values (in CGS units) appropriate for this object.

2.1 MAGNETIC FIELD IN THE HII SHELL

The magnetic field in the ambient medium is assumed to be weak enough to permit similarity solutions. This is certainly true as long as $B_0 \ll 8\pi c_i \sqrt{\rho_0} = 3.2\mathrm{mG}$, where $c_i = 10^6\mathrm{cm/s}$ is the isothermal speed of sound in the HII layer, ρ_0 is the density and B_0 the field strength in the ambient molecular cloud. The value for B_0 quoted above is certainly much larger than the average observed value which is between 10 and 100 μG. In the following we calculate B_i, the field in the HII region by assuming that a uniform (on a scale of parsecs) and parallel ambient magnetic field B_0, frozen into the circumstellar gas, is swept up into a thin shell. From similarity solutions we get that $R_b = 0.76(L_w/\rho_0)^{1/5} t^{3/5}$, where R_b is the radius of the bubble, $L_w = 10^{36}\mathrm{erg}$ is the wind mechanical luminosity and $\rho_i = 0.78\rho_0(\dot{R}_b/c_i)^2$.

The thickness of the partly ionized shell can be inferred from ionization equilibrium to be

$$l = 0.13 \frac{S_* m_a^2 c_i^4}{b \rho_0^2 R_b^2 \dot{R}_b^4} \tag{1}$$

and thus the fraction of the ionized mass is given by $f_i = 3\rho_i l/(\rho_0 R_b)$. Conservation of magnetic flux Φ at polar angle θ yields $\Phi = \pi B_0 R_b^2 \sin^2\theta$, in the shocked layer and therefore a fraction $f_i\Phi$ in the ionized part. The field strength will then obey

$$B_i = f_i \Phi/(2\pi R_b l \sin\theta) = 1.15 B_0 (\dot{R}_b^2/c_i^2) \sin\theta \, . \tag{2}$$

In this simple model the magnetic field decreases from its maximum value at the equatorial plane, i.e. perpendicular to the orientation of the large scale field, where $\theta = \pi/2$, to zero at the poles.

2.2 MAGNETIC FIELD IN THE STELLAR WIND

Consider an early type star, rotating at an angular velocity Ω_*; disregarding the complicated details of the stellar field topology, there will be lines of force leaving the northern and reentering the southern hemisphere. The field will be carried away by the wind and wound up into an Archimedian spiral with a neutral current sheet in the equatorial plane of the star. The field components in the radial and azimuthal direction in the unshocked stellar wind region are given by

$$B_r = B_*(R_*/r)^2 \,, \tag{3a}$$

$$B_\phi = B_*(\Omega_*/V_w)(R_*/r)^2(r - R_*)\sin\psi \,, \tag{3b}$$

where R_* and B_* are the stellar radius and surface field, V_w is the wind speed and ψ is the polar angle with respect to the axis of rotation. The wind will expand freely as long as the magnetic energy density is less than the kinetic energy density, meaning that $B_* \ll \sqrt{2\dot{M}_w V_w^3/(\Omega_* R_*^2)} = 2800G$ for typical values of $\Omega_* = 10^{-5}\mathrm{s}^{-1}$, $R_* = 10^{12}\mathrm{cm}$, $V_w = 2\times10^8\mathrm{cm/s}$ and a mass loss rate of $\dot{M}_w = 6.3\times10^{19}$ g/s. This condition will be satisfied in most cases.

Next we calculate the magnetic field in the hot bubble. We make the approximation that the shocked gas in the bubble is isothermal and the density therefore is uniform, given by $\rho_b = 3M_b/(4\pi R_b^3) = 0.54\dot{M}_w(\rho_0/L_w)^{3/5}t^{-4/5}$. Then the radial distance of fluid elements emitted at time $\tau < t$ is simply $R(\tau;t) = (1-\tau/t)^{1/3}R_b$. If the flow is not disturbed by magnetic stresses and the field is coupled to the streamlines of the gas, then a parcel of fluid emitted at $\tau + \delta\tau$ is offset in the radial direction by an amount $\delta R = (\partial R/\partial\tau)\delta\tau = -(R_b\delta\tau)/(3t(1-\tau/t)^{2/3})$. At polar angle ψ the transverse displacement due to the rotation of the underlying star is $-R\Omega_*\sin\psi\delta\tau$. While the radial component remains unchanged by the passage through the shock and is given by equation (3a), the azimuthal field is

$$B_\phi = 3\Omega_* B_* R_*^2 t \frac{R}{R_b^3}\sin\psi \,. \tag{4}$$

Taking $t = 2.2\times10^{11}$s as the age for NGC 6334(A), the condition that the energy density in the magnetic field must be less than in the thermal gas requires $B_* < 450$G. The field free layer in the equatorial plane arises by reconnexion of antiparallel field lines. We wish to estimate in the following under which conditions it is possible to reconnect a substantial fraction of the magnetic field in the hot bubble. The local Alfvén speed is $v_A = B/\sqrt{4\pi\rho_b} = 2.62\times(\Omega_* B_* R_*^2/\dot{M}_w^{1/2})(\rho_0/L_w)^{3/10}R\sin\psi t^{-2/5}$. The rate of reconnexion in the bubble is $\sigma = C(v_A/R_b)$, where the parameter C, containing the detailed physics, is constrained by the condition that if reconnexion is to become important we ought to have $\sigma t > 1$ or $C > R_b/(v_A t) = 2.2\times10^{-3}$ for the numbers given for NGC 6334(A) at the present age. The process of reconnexion is still not well understood; C can be as large as ~ 0.1 if Petschek's (1964) mechanism is at work and as low as $(R_m)^{-1/2}$, where R_m is the magnetic Reynolds number. Since the magnetic resistivity is fairly low in a hot, tenuous plasma, R_m is fairly large and therefore we expect instabilities to enhance the reconnexion rate considerably. Hence the plasma in the equatorial plane will be squeezed towards the contact surface at some fraction of the Alfvén speed and create a field free boundary layer, which allows free turbulent mixing with the HII gas.

3. Turbulent Mixing between the HII Region and the Hot Bubble

We have shown (Kahn and Breitschwerdt, 1989) that turbulent mixing driven by the acoustic instability occurs predominantly near the polar caps (below a critical angle θ) of the bubble, where the field in the HII gas is sufficiently weak. The rate of mass addition was found to be $\dot{\mathcal{M}} = 2.7 \times 10^{-5} L_w^{2/5} \rho_0^{8/5} c_i^3 B_0^{-2} t^{6/5}$. NGC 6334(A) will suffer then from considerable radiative losses at present due to mixing, if B_0 is less than 4×10^{-5}G.

4. Conclusions

We expect turbulent mixing to be important for most of the time during the dynamical evolution of a wind-blown HII region, even in the presence of magnetic fields. The exchange of mass and energy will cause significant changes in the flow in particular regions of local enhanced cooling, and could be responsible for high velocity flows and the presence of highly ionized species.

5. References

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DOGIEL: It seems to me that your suggestions – Kolmogorov's spectrum of a gas turbulence and magnetic field reconnections with the Alfvén velocity – are contradictory. The first assumption means, indeed, that magnetic fields are weak, but the second one is valid only for the cases of strong magnetic fields when the magnetic field energy density is larger than an energy density of gas motions. However, in the last case the turbulence spectrum is not Kolmogorov's one.

BREITSCHWERDT: We use Petschek's mechanism here to describe magnetic reconnection, where the rate is proportional to the Alfvén velocity over the characteristic length scale on which this process operates. The physics is hidden in a proportionality constant which itself depends on the local magnetic field strength. This has to be borne in mind if one deduces quantitative results from that. In addition, the velocities at all scales are very much subsonic so that we are allowed to calculate the reconnection rate in a quasi-stationary fashion.

C. NORMAN: Mordecai MacLow and myself have performed numerical simulations of thin shell instabilities of the sort studied by Vishniac. These instabilities are not exponential, but rather bistable. Our simulations confirm Vishniac's predictions and show that the shell buckles, but does not turn over and become turbulent. Could you comment on the nature of your thin shell instability, and how it differs from Vishniac's.

BREITSCHWERDT: The instability we refer to here is of an acoustic nature and arises from the fact that the ionization front must be well shielded from the stellar Lyc flux. Therefore disturbances travelling parallel to the HII/bubble interface induce gas motions perpendicular to the ionization front, which can do work on the waves and thereby amplify it. The amplitudes will then grow exponentially.

SPANGLER: Radio wave scattering observations can provide information on turbulence in some of these HII regions. We have observations of several radio sources viewed through the Cygnus OB1 association, which is a good example of an interstellar bubble. These observations indicate the presence of turbulence with a Kolmogorov spectrum from $\sim 10^8$ cm to 1 parsec.

BREITSCHWERDT: Here we have applied our model of turbulent mixing to the bubble of a single early-type star, but it should also hold for an OB association.