

THE STATE OF THE GAS AROUND YOUNG STELLAR GROUPS

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RESUMEN

Presentamos una breve descripción de la evolución de la presión del gas en regiones de formación estelar, desde que se forma la nube materna hasta que los vientos de las estrellas recién formadas presurizan la región. Se describen los procesos que destruyen la nube y la forma en que se auto-limita el número total de estrellas recién formadas. La alta eficiencia de la formación estelar en brotes nucleares, es debida a las altas presiones del gas. También se describe brevemente la evolución de vientos lentos y masivos en regiones de alta presión.

ABSTRACT

We present the main features in the evolution of the gas pressure in star forming regions, from the formation of the parental cloud to the moment when the region is pressurized by interacting stellar winds. The main processes for cloud destruction and the self-limiting properties of star formation are described. The high star forming efficiency in nuclear starbursts is a consequence of the high gas pressures. The evolution of slow winds in highly pressurized region is also sketched.

Key words: **GALAXIES: STARBURST — HYDRODYNAMICS — ISM: BUBBLES — ISM: STRUCTURE — STARS: FORMATION**

1. INTRODUCTION

The evolution of gaseous galaxies is affected by a large number of stochastic factors. They are open systems and the properties of their interstellar gas are regulated by internal sources of energy and interactions with other galaxies. There are a series of different mechanisms, for both isolated and interacting galaxies, that are able to accumulate large gas masses and create star-forming clouds in relatively short time scales. These include cloud-cloud collisions, gravitational, magnetic and thermal instabilities, gas flows in a bar potential, tidal interactions and direct galaxy collisions. Some of these mechanisms are discussed in this volume (Boeker et al.; Borne et al.; Dultzin-Hacyan; Elmegreen; Friedli & Martinet; Lamb et al.; Moss & Whittle). Any one, or a combination, of these processes could be operative in different locations and at different moments in the host galaxy, and star formation is the end product of a series of successive condensations of the interstellar medium. Once a cloud is formed, however, a distinction should be made between molecular and self-gravitating clouds (see Franco & Co. 1986). The criterion to form molecular clouds is high opacity in the UV, to prevent molecule photo-destruction and this is achieved at column densities above $N_{\tau} \sim 5 \times 10^{20} (Z/Z_{\odot})^{-1} \text{ cm}^{-2}$, where Z is the metallicity and $Z_{\odot}$ is the solar value. In contrast, self-gravity becomes dominant when the column density becomes larger than $N_{sg} \sim 5 \times 10^{20} (P/P_{\odot})^{1/2} \text{ cm}^{-2}$, where P is the interstellar pressure and $P_{\odot}$ is the value at the solar circle. These two values are similar at the location of the Sun, but $N_{\tau} < N_{sg}$ in the inner Galaxy and $N_{\tau} > N_{sg}$ in the outer parts of the Milky Way. This difference has important consequences and may explain the observed radial trends of molecular gas in spirals: it is easier to form molecular clouds in the internal, chemically evolved parts of spiral galaxies. In any case, the transformation of gas into stars is due to a gravitational collapse, and self-gravity defines the structure of the star forming clouds.

2. SELF-GRAVITY

The formation of stellar groups (or isolated stars, if any) occurs in the densest regions, the cores, of massive and self-gravitating clouds. In our Galaxy, the *average* densities for giant molecular cloud complexes is in the range 10^2 to 10^3 cm^{-3} , but the actual densities in the dense cores is several orders of magnitude above these

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values: close to about $\sim 10^6 \text{ cm}^{-3}$ (e.g., Bergin et al. 1996; see recent review by Walmsley 1995). Moreover, recent studies of young stellar objects suggest the existence of even denser gas, with values in excess of 10^8 cm^{-3} (Akeson et al. 1996). Thus, the parental clouds have complex clumpy (and filamentary) structures, with clump-interclump density ratios of about $\sim 10^2$, or more, and temperatures ranging between 10 and 10^2 K . In addition, the existence of large non-thermal velocities, of several km s^{-1} , and strong magnetic fields, ranging from tens of μG to tens of mG (see Myers & Goodman 1988 and references therein), indicate large *total* internal pressures, up to more than five orders of magnitude above the ISM pressure at the solar neighborhood (which is about $10^{-12} \text{ dyn cm}^{-2}$). A simple estimate for isothermal, spherically symmetric, clouds (with a central core of constant density ρ_c and radius r_c , and an external diffuse envelope with a density stratification $\rho = \rho_c(r/r_c)^{-2}$), indicates that self-gravity provides these large total pressure values (see García-Segura & Franco 1996). In hydrostatic equilibrium, the pressure difference between two positions located at radii r_1 and r_2 from the center of the core is given by $\Delta P = -\int_{r_1}^{r_2} \rho g_r dr$, where g_r is the gravitational acceleration in the radial direction. The total pressure at the core center is

$$P(0) = P_0 = \frac{2\pi G}{3} \rho_c^2 r_c^2 + P(r_c) = \frac{8}{5} P(r_c) \simeq 2 \times 10^{-7} n_6^2 r_{0.1}^2 \text{ dyn cm}^{-2}, \quad (1)$$

where G is the gravitational constant, $P(r_c)$ is the pressure at the core boundary $r = r_c$, $n_6 = n_c/10^6 \text{ cm}^{-3}$, and $r_{0.1} = r_c/0.1 \text{ pc}$. The corresponding core mass is

$$M_c \simeq \left(\frac{\pi P_0}{G} \right)^{1/2} r_c^2 \sim 10^2 P_7^{1/2} r_{0.1}^2 M_\odot, \quad (2)$$

where $P_7 = P_0/10^{-7} \text{ dyn cm}^{-2}$. For $P_7 \sim 1$ and a typical core size for galactic clouds, $r_{0.1} \sim 1$ (see Walmsley 1995) gives a value similar to the observationally derived core masses; in the range of 10 to $300 M_\odot$ (e.g., Snell et al. 1993). The pressure inside the core varies less than a factor of two between the center and $r = r_c$. Taking $r_{0.1} = 1$ and the maximum core density value, $n_c \sim 5 \times 10^6 \text{ cm}^{-3}$ (e.g., Bergin et al. 1996), the upper bound for the expected core pressures is about $P_0 \simeq 5 \times 10^{-6} \text{ dyn cm}^{-2}$. The large range in observed cloud properties obviously results in pressure fluctuations of a few orders of magnitude (both, from cloud to cloud and inside any given cloud), and it is meaningless to define an “average” cloud pressure value. Actually, given that star forming clouds have nested structures, in which dense fragments are embedded in more diffuse envelopes, different cloud locations have different total pressures. Also, the expected range of cloud pressures in our Galaxy should probably span from the ISM values, $P_7 \sim 10^{-5}$, at the very external cloud layers, up to $P_7 \sim 10$ inside the most massive star forming cores.

3. STELLAR RADIATION: H II REGIONS AND CLOUD DESTRUCTION

The initial structure and pressure of the gas in a star forming cloud is defined by self-gravity. Once young stars appear, the new energy input modifies the structure and evolution of the cloud. Low-mass stars provide a small energy rate and affect only small volumes, but their collective action may provide partial support against the collapse of their parental clouds, and could regulate some aspects of the cloud evolution (Norman & Silk 1980; Franco & Cox 1983; Franco 1984; McKee 1989; see also the paper by Bertoldi & McKee in this volume). In contrast, the strong radiation fields and fast stellar winds from massive stars are able to excite large gas masses and can even disrupt their parental clouds (e.g., Whitworth 1979; Franco et al. 1994). Also, they are probably responsible for both stimulating and shutting off the star formation process at different scales. The combined effects of supernovae, stellar winds, and H II region expansion destroy star-forming clouds and can produce, at some distance and later in time, the conditions for further star formation (e.g., Franco & Shore 1984; Palous et al. 1994). Thus, the transformation of gas into stars may be a self-limited and self-stimulated process (see reviews by Franco 1991, Ferrini 1992, and Shore & Ferrini 1994).

In the case of the dense star-forming cores, the sizes of either H II regions or wind-driven bubbles are severely reduced by the large ambient pressure (García-Segura & Franco 1996). In fact, the pressure equilibrium radii of ultra-compact H II regions are actually indistinguishable from those of ultra-compact wind-driven bubbles. When pressure equilibrium is reached, the UCHII radius is

$$R_{\text{UCHII,eq}} \approx 2.9 \times 10^{-2} F_{48}^{1/3} T_{\text{HII},4}^{2/3} P_7^{-2/3} \text{ pc}, \quad (3)$$

where F_{48} is the total number of ionizing photons per unit time in units of 10^{48} s^{-1} , and $T_{\text{HII},4} = T/10^4 \text{ K}$. For

the case of a strong wind evolving in a high-density molecular cloud core, the equilibrium radius of a radiative bubble is

$$R_{\text{WDB,eq}} = \left[\frac{\dot{M} v_{\infty}}{4 \pi P_0} \right]^{1/2} \simeq 2.3 \times 10^{-2} \left[\frac{\dot{M}_6 v_{\infty,8}}{P_7} \right]^{1/2} \text{ pc}, \quad (4)$$

where the mass loss rate is $\dot{M}_6 = \dot{M}/10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$, and the wind velocity is $v_{\infty,8} = v_{\infty}/10^8 \text{ cm s}^{-1}$. Thus, for dense cores with $r_c \sim 0.1 \text{ pc}$, the resulting UCHII and wind-driven bubbles can reach pressure equilibrium without breaking out of the core (i.e., they could be stable and long lived). Recently, Xie et al. (1996) have found evidence indicating that this is probably the case: the smaller UCHII seem to be embedded in the higher pressure cores.

If the limit to star formation is due to photoionization, the maximum number of OB stars is given by the number of H II regions required to completely ionize the core (Franco et al 1994), $N_{\text{OB}} = (1 - \epsilon)M_c/M_i$, where M_i is the ionized mass and ϵ is the star forming efficiency. This means that the maximum number of massive stars that can be formed within a core is

$$N_{\text{OB}} \approx 3 \frac{M_{c,2} n_6^{3/7}}{F_{48}^{5/7} (c_{s,15} t_{MS,7})^{6/7}} \quad (5)$$

where $M_{c,2}$ is the core mass in $10^2 \text{ M}_{\odot}$, $c_{s,15}$ is the H II region sound speed in units of 15 km s^{-1} , and $t_{MS,7}$ is the mean OB star main sequence lifetime in 10^7 yr . Clearly, for increasing core densities, the value of R_0 decreases and the resulting number of OB stars increases. For the gas in nuclear regions of galaxies, the population of clouds are denser and more compact due to the intrinsic larger ISM pressures in the inner regions. The corresponding star forming clouds should also be denser than in the rest of the disk, and a larger number of stars can be formed per unit mass of gas. Thus, *nuclear starbursts can be a natural consequence of the higher ISM pressure values* (an additional bursting star formation mode can also be associated to a delayed energy input, see Parravano 1996).

When stars are located near the edge of the core, and depending on the slope of the external density distribution, both H II regions and wind-driven bubbles can accelerate and flare out with a variety of hydro-dynamical phenomena. These include supersonic outflows, internal shocks, receding ionization fronts, fragmentation of the thin shell, etc (e.g., Tenorio-Tagle 1982; Franco et al. 1989, 1990; García-Segura & Mac Low 1995a,b). Thus, no static solution exists in this case and the pressure difference between the H II regions and the ambient medium begins to evaporate gas from the cloud. This represents a clear and simple physical mechanism for cloud destruction and, as the number of OB stars increases, more expanding H II regions form and limit the rate of new star formation by ionizing the surrounding molecular gas (Franco et al. 1994). Eventually, when the whole cloud is completely ionized, star formation ceases. The total cloud mass ionized by an average OB star, integrated over its main sequence lifetime, is

$$M_i(t) \approx \frac{2\pi}{3} R_0^3 \mu_p n_0 \left[\left(1 + \frac{5c_s t_{MS}}{2R_0} \right)^{6/5} - 1 \right] \quad (6)$$

where R_0 is the initial radius at the average cloud density, n_0 , μ_p is the mass per gas particle, c_s is the sound speed in the H II region, and t_{MS} is the main sequence lifetime of the average OB star. For a cloud of mass M_{GMC} , with only 10% of this mass in star-forming dense cores, the number of newly formed OB stars required to completely destroy it is

$$N_{\text{OB}} \sim 30 \frac{M_{\text{GMC},5} n_3^{1/5}}{F_{48}^{3/5} (c_{s,15} t_{MS,7})^{6/5}}, \quad (7)$$

where $M_{\text{GMC},5} = M_c/10^5 \text{ M}_{\odot}$, $n_3 = n_0/10^3 \text{ cm}^{-3}$, $c_{s,15} = c_s/15 \text{ km s}^{-1}$, and $t_{MS,7} = t_{MS}/10^7 \text{ yr}$. Assuming a standard IMF, this corresponds to a total star forming efficiency of about $\sim 5 \%$. For the average values of stellar ionization rates and giant molecular cloud parameters in our Galaxy, the overall star forming efficiency should be about 5%. Obviously, larger average densities and cloud masses can result in higher star formation efficiencies.

Summarizing, photoionization from OB stars can destroy the parental cloud in relatively short time scales, and defines the limiting number of newly formed stars. The fastest and most effective destruction mechanism is due to peripheral, blister, H II regions, and they can limit the star forming efficiency at galactic scales. Internal H II regions at high cloud pressures, on the other hand, result in large star forming efficiencies and they may be the main limiting mechanism in star forming bursts and at early galactic evolutionary stages (see Cox 1983).

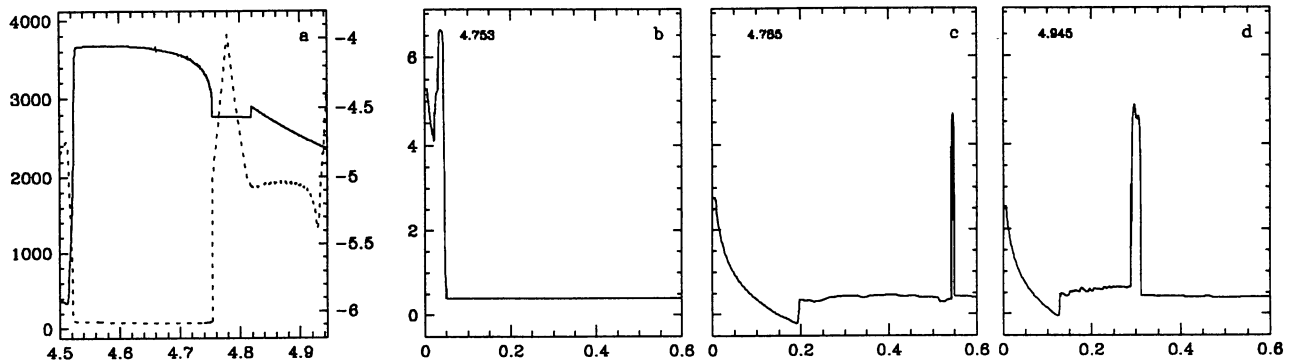


Fig. 1. Evolution of a wind-driven bubble in a high pressure medium. (a) Evolution of a stellar wind from a 35 M_⊙ star. Left scale: terminal wind velocity (km s⁻¹); right scale: mass loss rate (M_⊙ yr⁻¹); horizontal scale: time (millions of yr). (b)-(d): Evolution of the wind-driven bubble. The gas density (cm⁻³) is plotted in a logarithmic scale against the radial distance (pc). Evolutionary times (shown in the upper-left corner of each panel) are given in million years.

4. MECHANICAL ENERGY

As the cloud is dispersed, the average gas density decreases and the newly formed cluster becomes visible. The individual HII regions merge into a single photo-ionized structure and the whole cluster now powers an extended, low density, HII region. The stellar wind bubbles now can grow to larger sizes and some of them begin to interact. As more winds collide, the region gets pressurized by interacting winds and the general structure of the gas in the cluster is now defined by this mass and energy input (Franco et al. 1996).

Given a total number of massive stars in the cluster, N_{OB} , and their average mass input rate, $\langle \dot{M} \rangle$, the pressure due to interacting adiabatic winds is

$$P_i \sim \frac{N_{OB} \langle \dot{M} \rangle c_i}{4\pi r_{clus}^2} \sim 10^{-8} \frac{N_2 \langle \dot{M}_6 \rangle c_{2000}}{r_{pc}^2} \text{ dyn cm}^{-2}, \quad (8)$$

where $N_2 = N_{OB}/10^2$, $\langle \dot{M}_6 \rangle = \langle \dot{M} \rangle / 10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$, $r_{pc} = r_{clus}/1 \text{ pc}$ is the stellar group radius, and $c_{2000} = c_i/2000 \text{ km s}^{-1}$ is the sound speed in the interacting wind region. This is the central pressure driving the expansion of the resulting super-bubble before the supernova explosion stage. For modest stellar groups with relatively extended sizes, like most OB associations in our Galaxy, the resulting pressure is only slightly above the ISM pressure (i.e., for $N_2 \sim 0.5$ and $r_{pc} \sim 20$, the value is $P_i \sim 10^{-11} \text{ dyn cm}^{-2}$). For the case of rich and compact groups, as those generated in a starburst, the pressures can reach very large values. For instance, for the approximate cluster properties in starbursts described by Ho in this volume, $r_{pc} \sim 3$ and $N_2 > 10$, the resulting pressures can reach values of the order of $P_i \sim 10^{-7} \text{ dyn cm}^{-2}$, similar to those due to self-gravity in star forming cores. At these high pressures, the winds at the evolved red giant (or supergiant) phases cannot expand much and they reach pressure equilibrium at relatively small distances from the evolving star. Thus, the large mass ejected during the slow red giant wind phase is concentrated in a dense circumstellar shell. An example of this is shown in Figure 1, where the evolution of a wind-driven bubble around a 35 M_⊙ star is presented. Fig. 1a shows the wind velocity and mass-loss rate (dashed and solid lines, respectively: García-Segura, Langer, & Mac Low 1996). We ran the simulation only over the time spanning the red supergiant and Wolf-Rayet phases, and assume that the region is already pressurized by the main sequence winds from massive stars. We used the AMRA code, as described by Plewa & Różyczka in this volume. During the RSG phase, the wind-driven shell is located very close ($R \sim 0.04 \text{ pc}$) to the star due to a very low ram-pressure of the wind (Fig. 1b). Later on (Fig. 1c), the powerful WR wind pushes the shell away from the star to the maximum distance of $R \approx 0.54 \text{ pc}$. Still later, when the wind has variations, the shell adjusts its position accordingly, and reaches the distance $R \sim 0.3 \text{ pc}$ at the end of simulation (Fig. 1d). It must be stressed that the series of successive accelerations and decelerations of the shell motion during the WR phase will certainly drive flow instabilities and cause deviations from the sphericity assumed in our model. The role of these multidimensional instabilities in the evolution of the shell is currently under study (with 2-D and 3-D models), and the results will be presented in a future communication.

Regardless of the possible shell fragmentation, however, when the star explodes as a supernova, the ejecta will collide with a dense circumstellar shell. This interaction generates a bright and compact supernova remnant,

with a powerful photoionizing emission (i.e., Terlevich et al. 1992; Franco et al. 1993; Plewa & Różyczka this volume), that may also be a very strong radio source, like SN 1993J (see Marcaide et al. 1995). If the shell is fragmented, the ejecta-fragment interactions will occur during a series of different time intervals, leading to a natural variability in the emission at almost any wavelength (see Cid et al. 1996). This type of interaction is also currently under investigation, and further modeling will shed more light on the evolution of SN remnants in high-pressure environs.

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