Chapter: Dynamics of the ISM (4/15/13)

The interstellar gas is rarely static, motionless. When neutral gas is first ionized, its temperature shoots up to 10^4 K, and because $H^0 \rightarrow H^+ + e^-$ the particle number density doubles. The increasing pressure (P = nkT) drives the HII region to expand. Stars that have ionizing powers are also sources of fast stellar winds. The fast stellar wind dynamically interacts with the ambient medium to blow a bubble. Massive stars end their lives in supernova explosion, and the supernova ejecta further interacts dynamically with the ambient medium to form supernova remnants (SNRs).

Many objects have different names but share the same dynamical processes: planetary nebulae – fast wind sweeping up previous slow wind to form a bubble, interstellar bubble – fast wind sweeping up ambient ISM to form a bubble, circumstellar bubble – fast wind sweeping up previous slow wind to form a bubble, superbubble – large bubble blown by many stars collectively via winds and supernovae.



The global structure of the ISM consists of a mixture of phases: 10^6 K ionized, 10^4 K ionized, 10^2-10^3 K neutral atomic, and 10 K molecular components. The production and distribution of these different components involve dynamic processes as well.

Hydrodynamic Equations of Motion

Two reference systems: Lagrangian – following a fluid element Eulerian – fixed in space

Equation of Motion, or Momentum Equation

$$\rho \frac{D\vec{v}}{Dt} \equiv \rho \left[\frac{\partial \vec{v}}{\partial t} + \vec{v} \cdot \nabla \vec{v} \right] = -\nabla P - \frac{1}{8\pi} \nabla B^2 + \frac{1}{4\pi} \vec{B} \cdot \nabla \vec{B} - \rho \nabla \phi$$

In this equation D/Dt denotes the Lagrangian time derivative, and $\partial/\partial t$ denotes the Eulerian time derivative. ρ is the density, \vec{B} is the magnetic field, and ϕ is the gravitational potential.

Poisson's Equation

$$\nabla^2 \phi = 4\pi G \rho$$

The gravitational forces in HII regions, planetary nebulae and supernova remnants are usually negligible.

Equation of Continuity, or Mass Conservation

$$\frac{D\rho}{Dt} \equiv \frac{\partial\rho}{\partial t} + \vec{v}\cdot\nabla\rho = -\rho\nabla\cdot\vec{v}$$

Energy Equation

$$\frac{DU}{Dt} \equiv \frac{D}{Dt} \left(\frac{3}{2} \sum_{j} N_{j} kT \right) = G - L + \frac{P}{\rho} \frac{D\rho}{Dt} - U \nabla \cdot \vec{v}$$

where U is the internal kinetic energy per unit volume; $\frac{P}{\rho} \frac{D\rho}{Dt}$ is the heating from compression, and $U \nabla \cdot \vec{v}$ is the dilation effect.

Ionization Equation

$$\frac{DN(X^{+i})}{Dt} = -N(X^{+i}) \int_{\nu_i}^{\infty} \frac{4\pi J_{\nu}}{h\nu} a_{\nu}(X^{+i}) d\nu + N(X^{+i+1}) N_e \alpha_A(X^{+i}, T) + N(X^{+i-1}) \int_{\nu_{i-1}}^{\infty} \frac{4\pi J_{\nu}}{h\nu} a_{\nu}(X^{+i-1}) d\nu - N(X^{+i}) N_e \alpha_A(X^{+i-1}, T) - N(X^{+i}) \nabla \cdot \bar{v}$$

For an ideal gas, $P = \frac{\rho kT}{\mu}$, where μ is the mean mass per particle.

For an ideal gas undergoing an adiabatic process, $P \propto \rho^{\gamma}$, where $\gamma = C_P/C_V$, and $\gamma = 5/3$ for a monatomic gas, and $\gamma < 5/3$ for atoms or molecules with energy levels.

Pressure disturbances propagate as sound waves, and the source velocity c is given by $c^2 = dP/d\rho$.

$$\frac{dP}{d\rho} = \gamma \ \frac{P}{\rho} = \frac{\gamma kT}{\mu} = c^2$$

If the period of the sound wave is much larger than the cooling time t_T , the kinetic temperature will remain closely equal to T_E , the value in radiative equilibrium, hence

$$c^{2} = \frac{dP}{d\rho} = \frac{kT_{E}}{\mu} + \frac{\rho k}{\mu} \frac{dT_{E}}{d\rho} = \frac{kT_{E}}{\mu} \left(1 + \frac{\rho}{T_{E}} \frac{dT_{E}}{d\rho}\right)$$

In HII regions, $\frac{dT_E}{d\rho} \cong 0$ for $n_e < 10^2 \text{ cm}^{-3}$, hence $c^2 = \frac{kT_E}{\mu}$, corresponding to $\gamma = 1$. This is thus called the isothermal sound velocity.

Isothermal sound velocity $c \sim 10 \text{ km s}^{-1}$ at 10^4 K , and $\sim 100 \text{ km s}^{-1}$ at 10^6 K .

If a pressure disturbance has a large amplitude, the front of the pulse steepens because the sound velocity is higher in the compressed region, leading to a nearly discontinuous shock front.

If a pressure disturbance travels at speed faster than the sound velocity, a shock wave is formed. Supersonic motion \rightarrow shocks.

Shock Fronts

Assume a one-dimensional disturbance propagating through a homogeneous medium with a constant velocity u_1 . In the reference frame traveling with u_1 , the flow is steady, and all quantities are functions of x only.

SHOCK FRONTS



Figure 10.1 Schematic diagram of radiating shock. In the upper diagram the fluid is shown moving from left to right. At x_1 the fluid enters the shaded area, representing the nonradiative shock, followed by the transition region, where the temperature drops by radiation. The changes of density and temperature are indicated schematically in the lower figure.

Case 1: Perfect gas. $\vec{B} = 0$

Three "jump conditions":

- 1. Conservation of matter $\rho_1 u_1 = \rho_2 u_2$
- 2. Conservation of momentum $P_1 + \rho_1 u_1^2 = P_2 + \rho_2 u_2^2$
- 3. Conservation of energy $u_1^2 + \frac{2\gamma}{\gamma-1} \frac{P_1}{\rho_1} = u_2^2 + \frac{2\gamma}{\gamma-1} \frac{P_2}{\rho_2}$

If no radiation occurs, the rate of increase of fluid energy = the rate of work done on the fluid by pressure forces:

$$u_2(\frac{1}{2}\rho_2 u_2^2 + \mathcal{U}_2) - u_1(\frac{1}{2}\rho_1 u_1^2 + \mathcal{U}_1) = u_1 P_1 - u_2 P_2$$

where \mathcal{U} is the internal energy of the fluid per unit volume.

For a perfect gas, $\mathcal{U} = \frac{1}{\gamma - 1} P$, thus

$$u_1^2 + \frac{2\gamma}{\gamma - 1} \frac{P_1}{\rho_1} = u_2^2 + \frac{2\gamma}{\gamma - 1} \frac{P_2}{\rho_2}$$

The three jump equations can be solved:

$$\frac{P_2}{P_1} = \frac{2\gamma}{\gamma+1} \mathcal{M}^2 - \frac{\gamma-1}{\gamma+1}$$
$$\frac{u_2}{u_1} = \frac{\rho_1}{\rho_2} = \frac{\gamma-1}{\gamma+1} + \frac{2}{\gamma+1} \frac{1}{\mathcal{M}^2}$$

where \mathcal{M} is the "Mach number", and

$$\mathcal{M}^2 = \frac{\rho_1 \ u_1^2}{\gamma \ P_1} = \frac{u_1^2}{c_1^2}$$

For large \mathcal{M} ,

$$\frac{\rho_2}{\rho_1} = \frac{\gamma + 1}{\gamma - 1}$$
 and $P_2 = \frac{2 \ \rho_1 \ u_1^2}{\gamma + 1}$

For an adiabatic shock, $\gamma = 5/3$, and $\rho_2/\rho_1 = 4$. **** Adiabatic shocks give a density increase if a factor of 4.

If radiative energy loss is efficient, the cooling behind the shock front leads to compression. If $\gamma = 1$, $\rho_2/\rho_1 = \mathcal{M}^2 = u_1^2/c^2$. A large compression is possible. **** Isothermal shocks can produce large density increases.

Case 2. Hydromagnetic Shocks

If a magnetic field parallel to the shock front is present, the magnetic pressure needs to be added in the momentum equation.

$$P_1 + \rho_1 u_1^2 + \frac{B_1^2}{8\pi} = P_2 + \rho_2 u_2^2 + \frac{B_2^2}{8\pi}$$

In a plasma, the magnetic field lines are "frozen" into the fluid:

$$\frac{B_1}{\rho_1} = \frac{B_2}{\rho_2}$$

In a strong adiabatic shock, density increases by a factor of 4, the magnetic pressure increases by a factor of 16. If $B_1 \leq 3 \times 10^{-6}$ G, $B_2^2/8\pi$ is less than 1/5 of P_2 for typical values of ρ_1 and u_1 in the ISM. Such a weak magnetic field does not affect the shock front significantly.

For an isothermal shock, $P = \rho c^2$. If the Alfvén velocity, $V_A = \sqrt{B^2/4\pi\rho}$, ahead of the shock is much larger than the isothermal sound velocity c, then $\rho_2/\rho_1 = 2^{1/2}u_1/V_{A1}$. The criteria for a strong shock now requires u_1 to be large compared to V_{A1} , instead of c.

If $B = 3 \times 10^{-6}$ G, ρ_1 in an HI cloud is 4.7×10^{-23} g cm⁻³, $V_A = 1.2$ km s⁻¹ $\sim c$. The compression is much less than that for an isothermal shock without magnetic field.

Observations of Shocks (Practical Questions)

What is the post-shock temperature?

Isothermal Shock - post-shock temperature same as pre-shock temperature. Adiabatic Shock -

$$kT = \frac{P_2}{\rho_2/\mu} = \frac{2\rho_1 u_1^2}{\gamma + 1} \frac{\mu}{\rho_2} = \frac{2\mu u_1^2}{\gamma + 1} \frac{\rho_1}{\rho_2}$$

A strong adiabatic shock has $\rho_2/\rho_1 = 4$ and $\gamma = 5/3$; therefore,

$$kT = \frac{3}{16}\mu u_1^2$$

In an ionized medium, $\mu \approx 0.5 \ m_H = 8.37 \times 10^{-25} \ g$ $u_1 = 300 \ \text{km s}^{-1} \rightarrow T = 1 \times 10^6 \ \text{K}$ (e.g. SNRs) $u_1 = 3,000 \ \text{km s}^{-1} \rightarrow T = 1 \times 10^8 \ \text{K}$ (e.g. fast stellar winds)

Shock Velocity vs Shell Expansion Velocity

For adiabatic shocks, $\rho_2/\rho_1 = 4$, and $u_1/u_2 = 4$. The observed emission comes from the post-shock region, where material flows downstream at 1/4 of the shock velocity u_1 . Therefore, the observed expansion velocity $= u_1 - u_2 = \frac{3}{4}u_1$.

Shock velocity $(u_1) = \frac{4}{3} \times$ observed expansion velocity.

Stellar Wind Blown Bubbles

Fast stellar winds sweep up the ambient medium to form bubbles. Two theories have been proposed to explain bubbles.

(1) Energy-conserving bubbles (Weaver et al. 1977, ApJ, 218, 377)

In this model, the shocked stellar wind does not cool; the hot, shocked stellar wind forms a layer of hot gas whose pressure drives the expansion of the swept-up medium. See Figures below.



(2) Momentum-conserving bubbles (Steigman, et al. 1975, ApJ, 198, 575) In this model, shocked stellar wind cools rapidly, and reaches the inner wall of the shocked interstellar gas shell. That is, the region (b) in the above figure has zero thickness.

For energy-conserving bubbles,

$$r(t) = \left(\frac{25\dot{M}V_w^2}{12\pi\rho_0}\right)^{1/5} t^{3/5}$$

For momentum-conserving bubbles,

$$r(t) = \left(\frac{3\dot{M}V_w}{2\pi\rho_0}\right)^{1/4} t^{1/2}$$

where r is the radius, \dot{M} is the stellar wind mass loss rate, V_w is the terminal stellar wind velocity, ρ_0 is the ambient density, and t is the time.

Bubble Observations vs Theories

According to theory, a massive star with strong stellar wind in a dense medium will blow a bubble; however, not every O star is seen in a bubble. To solve this puzzle, N11B has been selected to search for bubbles around O stars. N11B is an HII region ionized by a young OB association that hosts O3 stars. Hubble Space Telescope image of N11B (shown below) does not show obvious bubbles.



High-dispersion spectra were obtained of regions around O stars in N11B. It is then found that O stars are indeed inside expanding shells, the presumed bubbles. The expansion velocities of these shells are $15-25 \text{ km s}^{-1}$, not too much higher than the sound velocity of 10 km s⁻¹. The outer shocks of these bubbles are not very strong, thus no strong compression is expected to produce clear shell morphology.

X-ray observations of wind-blown bubbles show that both the hot gas temperature and X-ray luminosity are lower than expected from theories. It is likely that the ambient medium is clumpy and the evaporation and ablation of the clumps "poison" the hot interior, lowering the temperature and X-ray emissivity.

The interface between hot interior and the cold shell has been detected in O VI and N V absorption lines. Few high-quality observations exist.

Detailed observations of size, expansion velocities, and stellar winds of superbubbles also show discrepancy from models. Observations of bubbles and superbubbles show expansion velocities slower than those expected from the observed stellar wind input and shell size. Superbubble ages implied from the size and expansion velocities are also smaller than the stellar ages of the central OB associations. There are still many puzzles about bubbles and superbubbles.

Supernova Remnants

Supernova (SN) ejecta interacts with the ambient circumstellar/interstellar medium. A shock will begin to form when the ejecta has swept over a distance comparable to the mean free path. Protons moving through a neutral HI medium with N(HI) = 1.2cm⁻³ at a velocity of 20,000 km s⁻¹ have a mean free path of ~500 pc, and the mean travel time is 2.5×10^4 yr. If this is true, SN ejecta can expand out of the galactic disk without forming any shocks. Magnetic field must be taken into account. For a proton moving across a 3 μ G magnetic field at 20,000 km s⁻¹, its gyration radius is only 10¹¹ cm (~ 3 × 10⁻⁸ pc). Thus, a hydromagnetic shock will form.

The evolution of a supernova remnant (SNR) is divided theoretically (simplistically) into four stages:

- (1) free expansion;
- (2) Sedov phase, when swept-up mass \sim ejecta mass, adiabatic shock;
- (3) snowplow phase, when cooling is important; isothermal shock; expansion is driven by momentum;
- (4) dissipated into the ISM.

It is now realized that "free expansion" is expanding into the circumstellar material ejected by the star before the SN explosion. Roger Chevalier was the first one taking into account the circumstellar material. The first stage is thus called "Chevalier phase" sometimes by SNR researchers.

Sedov 1959, Similarity and Dimensional Methods in Mechanics

A point blast with adiabatic expansion, $\gamma = 5/3$ for a perfect gas, the similarity solution is:

$$R_s = 12.8 \left(\frac{t}{10^4}\right)^{2/5} \left(\frac{E_{51}}{n_0}\right)^{1/5} \text{pc}$$
$$v_s = 500 \left(\frac{t}{10^4}\right)^{-3/5} \left(\frac{E_{51}}{n_0}\right)^{1/5} \text{km s}^{-1}$$

$$T_s = 3.4 \times 10^6 \left(\frac{t}{10^4}\right)^{-6/5} \left(\frac{E_{51}}{n_0}\right)^{2/5} \mathrm{K}$$

where t is the age in units of year, E_{51} is the explosion energy in units of 10^{51} ergs (i.e. 1 foe).

During the Sedov phase, the kinetic energy of the SNR shell is $\sim 30\%$ the initial explosion energy E_0 . The shell density is $n_s = 4n_0$.

During the snowplow phase, the expansion is initially driven by the pressure of the SNR interior, and finally by just the momentum of the shell. Toward the end of the snowplow phase, the shell kinetic energy is only 3-4% of E_0 .

