Advanced Cosmology Course: ISM and Star Formation

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June, 9 2015

Outline

- Main components of the ISM
- Molecules in the Cold Neutral Medium
- Giant Molecular Clouds
- Kennicutt-Schmidt relation
- Application: evolution of a Supernova Remnant

The InterStellar Medium

The space between stars in a galaxy is referred to as the *InterStellar Medium* or **ISM** and it includes

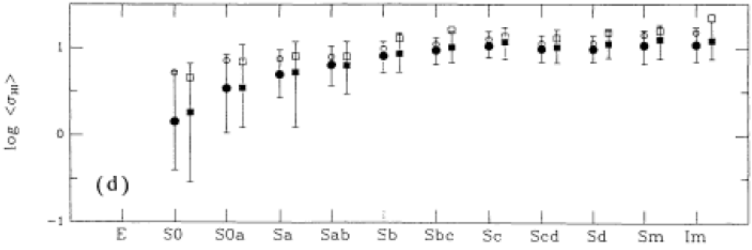
- atomic gas
- ionized gas
- molecular gas
- dust
- cosmic rays

In the Milky Way the ISM makes up 10-15% of the visible mass and 99% of it is gas, while only 1% is dust.

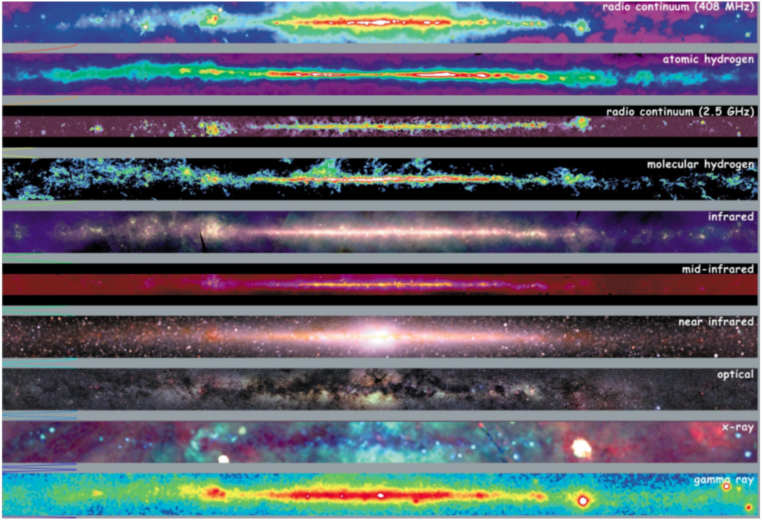
The gas mass and fraction vary systematically along the Hubble sequence

Early-type galaxies typically have very little gas (i.e. $M_{\rm gas} < 10^7$ M$_\odot$) and very little dust

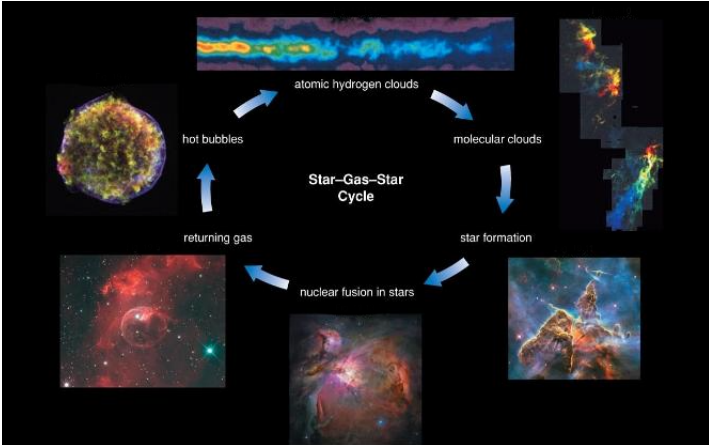
Late-type galaxies contain the majority of gas and dust, with masses up to several 10^{10} M$_\odot$



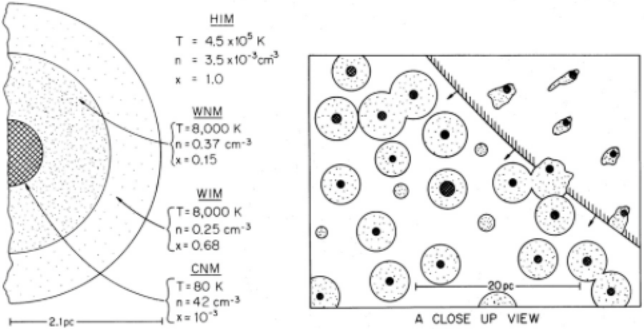
The InterStellar Medium



The InterStellar Medium



3-phase ISM (see McKee & Ostriker 1977)



3-phase model

- **Cold Neutral Medium (CNM)**, $T \sim 30 - 100$ K, $n_{\rm H} > 10$ cm^{-3}
  Dense, atomic (HI) and molecular (H_2, CO, ...) species, dust clouds.
  Only a few percents of the total volume but it contains most of the mass of the ISM.
  Confined to thin disk. It's the place where stars are born.
- **Warm Neutral Medium (WNM) / Warm Ionized Medium (WIM)**, $T \sim 10^3 - 10^4$ K, $n_{\rm H} \sim 1$ cm^{-3}
  Neutral and ionized gas, energized by UV radiation from stars, most of the total ISM volume in disk galaxies
- **Hot Ionized Medium (HIM)**
  $T \sim 10^5 - 10^6$ K, $n_{\rm H} \ll 1$ cm^{-3}
  Galactic corona
  Almost fully ionized, energized mainly by SN shocks

Thermal instability

Thermal instability can give an explanation for the different phases of the ISM.

We can assume that the ISM is in thermal equilibrium and ask ourselves what happens if we perturb that state.
From the equation of the conservation of energy we have

$$\frac{\partial \varepsilon}{\partial t} + \nabla \cdot \left[(\varepsilon + P)\, \vec{v} \right] = \rho\, \dot{Q} \ \ ,$$

where ε is the energy density, P the pressure, $\vec{v}$ the velocity, ρ the density and Q the energy per unit mass.

Because one has

$$\mathrm{d}Q = T\, \mathrm{d}s$$

we have

Thermal instability

We can write the *heat equation**

$$\rho\,T\,\frac{\mathrm{D}s}{\mathrm{D}t} = -\rho\,\left[\rho\,\Lambda(\rho,T) - \Gamma(T)\right] \quad ,$$

where $\Lambda(\rho,T)$ and $\Gamma(T)$ are the cooling and heating function, respectively.

Let's define $\mathcal{L}$ to be

$$\mathcal{L} = \rho\,\Lambda(\rho,T) - \Gamma(T) \quad ,$$

so that if $\mathcal{L} > 0$ cooling dominates over heating.

Thermal instability

The equations of fluid-dynamics in absence of gravity and viscosity are

$$\begin{cases} \dfrac{\partial \rho}{\partial t} + \nabla \cdot (\rho \vec{v}) = 0 \\ \\ \rho \dfrac{\partial \vec{v}}{\partial t} + \rho \vec{v} \cdot \nabla \vec{v} = -\nabla P \\ \\ T \dfrac{\mathrm{D}s}{\mathrm{D}t} = -\mathcal{L} \end{cases}$$

If the gas is a polytrope, i.e. $P = K\rho^{\gamma}$, it can be shown that

$$T\,\mathrm{D}s = \frac{P}{\rho\,(\gamma - 1)}\,\frac{\mathrm{D}K}{K}$$

Thermal instability

Let's consider a small perturbation of a static equilibrium in thermal balance

$$\begin{cases} \vec{v} \rightarrow \vec{v}_0 + \delta\vec{v} = \delta\vec{v} \\ \mathcal{L} \rightarrow \mathcal{L}_0 + \delta\mathcal{L} = \delta\mathcal{L} \\ P \rightarrow P_0 + \delta P \\ K \rightarrow K_0 + \delta K \\ \rho \rightarrow \rho_0 + \delta\rho \end{cases}$$

where $|\delta\vec{v}| \ll 1$, $\delta\mathcal{L} \ll 1$, $\delta P \ll 1$, $\delta K \ll 1$, $\delta\rho \ll 1$.

Thermal instability

The linearized equation to solve re

$$\begin{cases} \dfrac{\partial\,\delta\rho}{\partial\,t} + \rho_0\,\nabla\cdot(\delta\vec{v}) = 0 \\[2ex] \rho_0\dfrac{\partial\,\delta\vec{v}}{\partial\,t} = -\nabla\delta P \\[2ex] \dfrac{\mathrm{D}\delta K}{\mathrm{D}t} = -\dfrac{\gamma-1}{\rho_0^{\gamma-1}}\delta\mathcal{L} \end{cases}$$

Thermal instability

Because $\delta\mathcal{L} = \left.\frac{\partial\mathcal{L}}{\partial P}\right|_\rho \delta P + \left.\frac{\partial\mathcal{L}}{\partial\rho}\right|_P \delta\rho$, we can write

$\frac{\mathrm{D}\delta K}{\mathrm{D}t} = -A^*\delta P - B^*\delta\rho$, where

$A^* = \frac{\gamma-1}{\rho_0^{\gamma-1}} \left.\frac{\partial\mathcal{L}}{\partial P}\right|_\rho$, $B^* = \frac{\gamma-1}{\rho_0^{\gamma-1}} \left.\frac{\partial\mathcal{L}}{\partial\rho}\right|_P$

Also, because we are dealing with a polytrope, we have

$$\mathrm{d}P = \rho^\gamma \mathrm{d}K + c_s^2 \mathrm{d}\rho$$

If we seek for solutions in the form

$$\delta\rho = \rho_1\, e^{i\vec{k}\cdot\vec{x}+\omega t}$$

we have

$$\begin{cases} \omega\rho_1 + \rho_0\, i\vec{k}\cdot\vec{v}_1 = 0 \\ \omega\rho_0\vec{v}_1 = -i\,\vec{k}\,P_1 \\ \omega K_1 = -A^* P_1 - B^*\rho_1 \\ P_1 = \rho_0^\gamma\, K_1 + c_s^2\rho_1 \end{cases}$$

After some algebra we can get the dispersion relation

$$E(\omega) = \omega^3 + A^*\omega^2\rho_0^\gamma + k^2 c_s^3 \omega - B^* k^2 \rho_0^\gamma$$

The solution is unstable if $\omega > 0$.

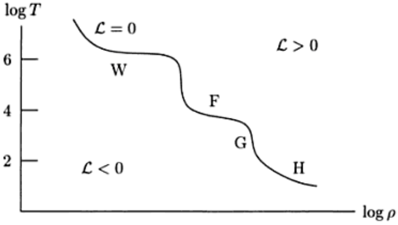
$E(\omega)$ has positive solution if $A^* < 0$ or $B^* > 0$. What does that mean?

$$\left.\frac{\partial \mathcal{L}}{\partial P}\right|_\rho < 0 \rightarrow \left.\frac{\partial \mathcal{L}}{\partial T}\right|_\rho < 0 \;\; \text{Parker condition}$$

$$\left.\frac{\partial \mathcal{L}}{\partial \rho}\right|_P > 0 \rightarrow \left.\frac{\partial \mathcal{L}}{\partial T}\right|_P < 0 \;\; \text{Field condition}$$

Thermal instability

Multi-phase equilibria are possible only if the gas can be thermally unstable and if there is only one T at which heating balances cooling, then (for a given pressure) only one phase is possible.



The H_2 molecule

H_2 is the most abundant molecule in the Universe and it mainly resides in Giant Molecular Clouds (GMCs)

H_2 is often involved as a *necessary* step for the formation of other molecules species and it is the main collisional partner for gas-phase reactions inside GMCs.

Detection of H_2

Although H_2 is very abundant, it is extremely difficult to detect. Let's understand why.

For a given electronic state q, a number of different rotational and vibrational states (with associated rotational quantum number J and vibrational quantum number v) can exist and the total vibration-rotation energy can be expressed as

$$E_q(v,J) = V_q(r_0) + h\nu_0\left(v + \frac{1}{2}\right) + B_v\, J(J+1)\ ,$$

where $V_q(r_0)$ is the minimum potential energy of the electronic state q and r_0 is the internuclear separation.

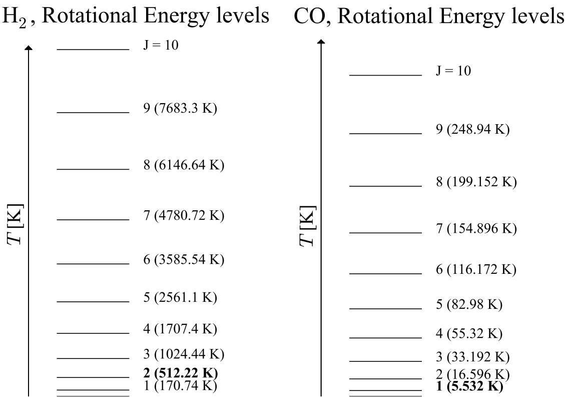
Detection of H_2

$$h\nu_0 = \hbar\sqrt{\frac{k}{\mu}} \;, B_v = \frac{\hbar^2}{2\mu r_0^2} \;,$$

μ is the molecule reduced mass and k is know as the force constant.

Because H_2 is the least massive molecule its energy levels are *more largely separated* with respect to any other molecule. Moreover, because it is a homonuclear molecular its does not have a permanent electric dipole and only *quadrupole* transitions are permitted.

Detection of H_2



Detection of H_2

The first rotational level accessible is more than **500 K** above the ground level.

H_2 mostly resides in the CNM, within GMCs, with $T \sim 10 - 30$ K!

Direct observation of H_2 is possible, but only in *rare* conditions:

- in the vicinity of a hot stars
- in post-shock gas

Observations are only possible with space-based telescopes because Earth's atmosphere is opaque to FUV radiation.

The H_2 chemical network: Three-body reactions

Dominant process at high gas density, i.e. $n \gtrsim 10^8$ cm^{-3}, and the dominant reaction is

$$\mathrm{H} \quad + \quad \mathrm{H} \quad + \quad \mathrm{H} \quad \rightarrow \quad \mathrm{H_2} \quad + \quad \mathrm{H} \ .$$

Even though its reaction rate is very small (i.e. $\sim 10^{-32}\,\mathrm{cm}^{-3}\,\mathrm{s}^{-1}$ at $T = 300$ K), gas cooling increases as soon as a significant fraction of H_2 is formed. As a result, gravitational collapse is further promoted, hence three-body reactions initiate a *cascade process* that almost leads to a complete conversion of hydrogen to its molecular form at densities as high as 10^{12} cm^{-3} (Palla et al. 1983).

Three-body reactions play a key role in the formation of protoplanetary discs, and also impact on the accretion physics onto Pop III protostars (Abel et al. 2002, Turk et al. 2011)

The H_2 chemical network: Gas-phase reactions

The rate coefficient of the direct association of two H atoms

$$H + H \rightarrow H_2 + \gamma$$

is as low as $\sim 10^{-30}$ cm^3 s^{-1}.

Gas-phase reactions are dominated by fast radiative association reactions with ions, through the H^- and the H_2^+ routes (see Saslaw et al. 67, Abel et al. 97):

$$\begin{array}{llllll} H & + & e^- & \rightarrow & H^- & + & \gamma \ , \\ H^- & + & H & \rightarrow & H_2 & + & e^- \ , \\ H^+ & + & H & \rightarrow & H_2^+ & + & \gamma \ , \\ H_2^+ & + & H & \rightarrow & H_2 & + & H^+ \ . \end{array}$$

The H_2 chemical network: Gas-phase reactions

At very high-redshift (i.e. $z \gtrsim 100$), the temperature of the cosmic microwave background (CMB) is large enough that the photo-detachment reaction

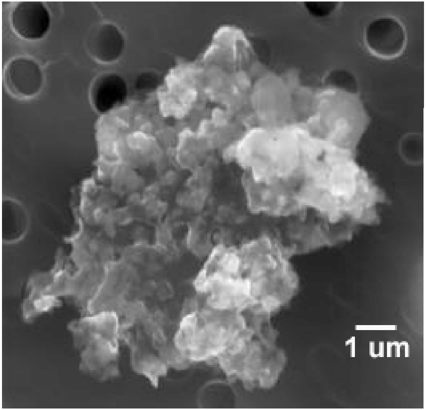
$$H^- + \gamma_{CMB} \rightarrow H + e^- ,$$

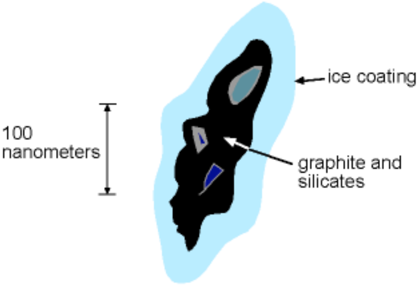
is highly favored and the abundance of H^- is considerably reduced, thus H_2 formation is dominated by the H_2^+ channel.

However, the overall contribution of H_2 through gas-phase reaction is minimal. For instance, Glover 2003 estimated that they can only account for a H_2 fraction of most equal to 10^{-3}.

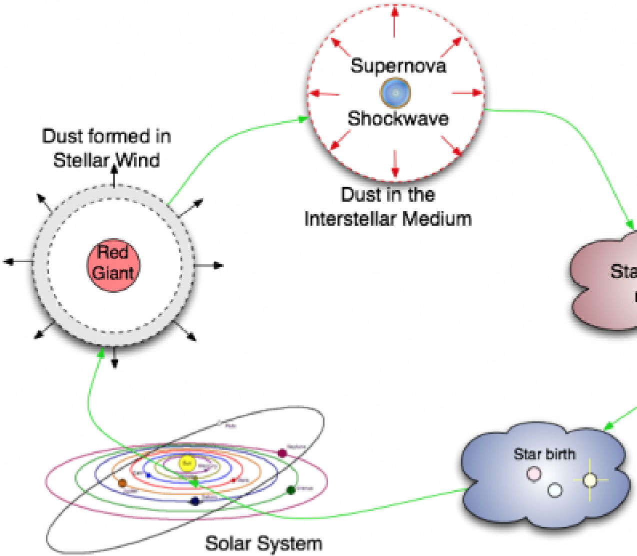
The dominant formation channel is through **dust grain reactions**

What are dust grains?





The H_2 chemical network: Dust grains reactions



The H_2 chemical network: Dust grains reactions

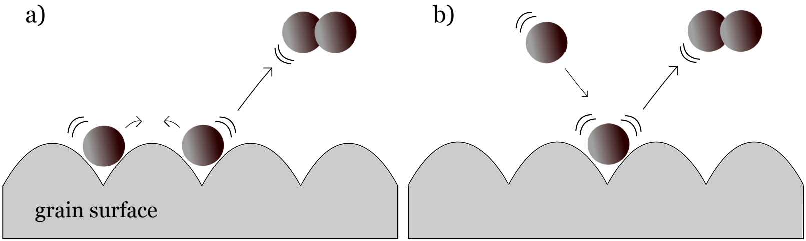
They consists of a **core** (of size $\sim 0.01\,\mu$m), which is mainly made up of silicates, iron and carbon.
During their life-cycle, they accrete atoms (mostly H, O, C and N) and form a mantel of ices of water, ammonia, carbon dioxide and methane around the core. Furthermore, interactions of their outer surface with UV radiation can generate complex molecular compounds in their **mantle**.

They make up as little as **1%** in mass in our Galaxy, but they presence is extremely important for the formation of molecules, because they act as *catalysts*.

The H_2 chemical network: Dust grains reactions

If sufficiently close to the grain surface, an atom can experience weak interactions (*van der Waals* forces) and enter either a weakly-bounded *physisorbed* site or a more strongly bounded *chemisorbed* site. *Adatoms* (i.e. atoms that have been absorbed to a surface) can move across the surface through thermal hopping (at high gas temperatures) or quantum tunneling (at low temperatures) and visit several sites.

Reactions between two or more adatoms known as: (a) *Langmuir-Hinshelwood kinetics*; or (b) reactions between adatoms and atoms from the gas phases that directly arrive into an occupied site known as *Eley-Rideal kinetics*.



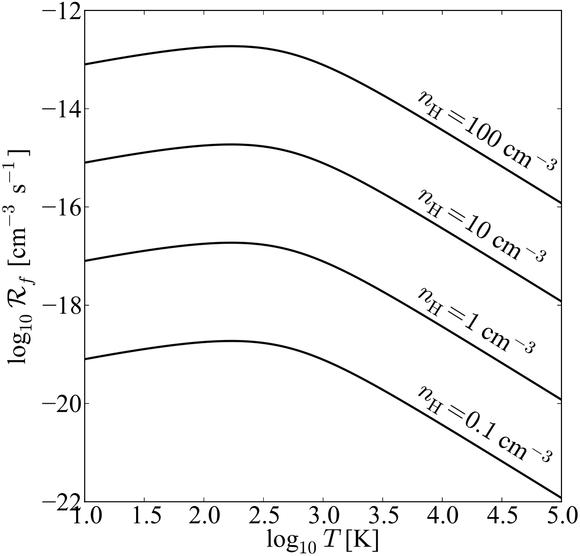
The H_2 chemical network: Dust grains reactions

The H_2 formation rate on dust grains can be parametrized as follows:

$$\mathcal{R}_f = \frac{1}{2}\, n_{\mathrm{HI}}\, v_{\mathrm{HI}}(T)\, n_d\, \sigma_g\, \varepsilon_{\mathrm{H_2}}(T_d)\, S_{\mathrm{H}}(T, T_d) \ \ ,$$

where the factor 1/2 takes into account that two HI atoms are necessary to form an H_2 molecule. Here n_{HI} is the number density of hydrogen atoms and $v_{\mathrm{HI}} = \sqrt{(8k_{\mathrm{B}}T)/(\pi m_{\mathrm{HI}})}$ is their mean thermal velocity at gas temperature T. The number density of dust is given by n_d, while σ_g is the grain cross sectional area and S_{H} is the sticking function (see Cazaux et al. 2004).

The H_2 chemical network: Dust grains reactions



The H_2 chemical network: Destruction processes

The dominant processes of H_2 destruction in the gas phase are collisional dissociations by electrons and atomic or molecular hydrogen:

$$\begin{array}{ccccccccc} H_2 & + & e^- & \rightarrow & H & + & H & + & e^- \;, \\ H_2 & + & H & \rightarrow & H & + & H & + & H \;, \\ H_2 & + & H_2 & \rightarrow & H_2 & + & H & + & H \;. \end{array}$$

They are important destruction mechanism only in the WNM, at $T \gtrsim 5000$ K.

The H_2 chemical network: Photodissociation

Photo-dissociation of molecular hydrogen is a two-step process, often referred to as the *Solomon process*, and it consists of the absorption of a UV photon, $\gamma_{\rm LW}$, which results in an excited molecular state, H_2^*, that leads to a decay in lower vibrational states (flourescent emission) or in the vibrational continuum of the ground state (photo-dissociation):

$$\mathrm{H_2} \quad + \quad \gamma_{\rm LW} \quad \rightarrow \quad \mathrm{H_2^*} \quad \rightarrow \quad \mathrm{H} \quad + \quad \mathrm{H} \ .$$

In the ground electronic state, H_2 absorbs radiation in two densely packed series of lines (the Lyman band – characterized by photon energies $E > 11.2$ eV or wavelengths $\lambda < 1108$ Å – and the Werner band – $E > 12.3$ eV, $\lambda < 1008$ Å).

The H_2 chemical network: Photodissociation

Radiative decay from the excited states leads to dissociation in approximately 15 per cent of the cases. Direct photo-dissociation would require photons with energy $E > 14.7$ eV, but these are principally absorbed by hydrogen atoms as they lie above the hydrogen photo-ionization threshold (13.6 eV, 912 Å). As a result, only photons between 912 Å $< \lambda <$ 1108 Å, the *Lyman-Werner* (LW) band, can photo-dissociate molecular hydrogen

The H_2 chemical network: Photodissociation

The rate of photodissiciation is estimated to be $\kappa = 4.2 \times 10^{-11}$ s^{-1} (Draine et al. 1996) for a unitary UV field in Habing units. This yields a very rapid photo-dissociation time scale, of order of 10^3 yr, however, it does not take into account shielding effects. In fact, H_2 can *self-shield*, meaning that, at the edge of a cloud, it effectively absorbs all available UV photons, and the molecules lying deeper do not receive any, thus the photo-excitation transition becomes optically thick. In addition, also the presence of dust can contribute to shielding from UV radiation and all of these effects should, in principle, be taken into account.

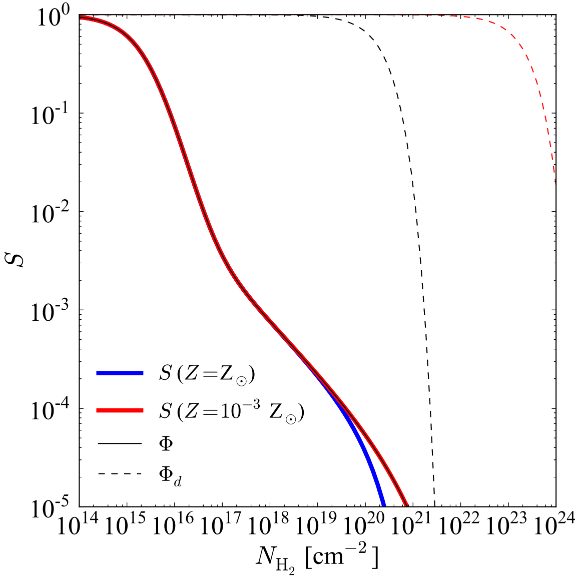
The H_2 chemical network: Photodissociation

It is conventional to introduce the shielding function S

$$S = \Phi \, \Phi_d \ ,$$

that ranges from 0 to 1 and describes the attenuation of UV radiation due to self-shielding (Φ) and dust absorption (Φ_d).

The H_2 chemical network: Photodissociation



Giant Molecular Clouds

Giant Molecular Clouds (GMCs) are gravitationally bound structure, shielded against UV radiation from stars that harbors most of the H_2 mass in a galaxy and host protostar in their inner parts.

Classification of clouds

Following Draine 2011, clouds can be classified according to their *visual extinction* A_V.

- If $A_V \lesssim 5$ coulds are transparent to most UV radiation and are referred to as *diffuse* or *translucent* clouds
- Clouds more opaque to radiation, with $A_V \sim 10$, are dubbed as *dark clouds*
- If they are opaque even at 8 μm ($A_V \gtrsim 100$) they are classified as *infrared dark clouds* (IRDC)

The main distinction between *Giant Molecular Clouds* (GMC) and *dark clouds* is in their mass, with the first ranging from 10^3 to 10^6 $M_\odot$ and the latter having masses $\lesssim 500$ $M_\odot$.

Classification of clouds

	GMC	Dark cloud	Star-forming clump	Protostellar cores
Size (pc)	$2 - 20$	$4 - 25$	$0.1 - 2$	$\lesssim 0.1$
Mass ($M_\odot$)	$10^3 - 10^6$	$10^2 - 10^4$	$10 - 10^3$	$\lesssim 10^2$
$\langle n_{H_2} \rangle$ (cm^{-3})	$10^2 - 10^4$	$10^2 - 10^4$	$10^4 - 10^5$	$10^4 - 10^6$
$\langle T \rangle$ (K)	$10 - 30$	$10 - 30$	$10 - 20$	$7 - 12$
A_V (mag)	$9 - 25$	$3 - 15$	$4 - 90$	$30 - 200$

Larsons relations: Velocity dispersion vs Size

The first one states that there is a correlation between the cloud size, L, and its velocity dispersion, σ_v:

$$\sigma_v \simeq 1.1 \left(\frac{L}{1\,\mathrm{pc}} \right)^{\gamma_L} \ \mathrm{km\,s^{-1}} \ .$$

The value of the exponent γ_L has long been debated and most recent measurements give $\gamma_L \sim 0.5 - 0.59$ (in disagreement with the value originally found by Larson, $\gamma_L = 0.38$) (Solomon et al. 87, Heyer et al. 04).

Larson second relation correlates σ_v and the cloud mass, M:

$$\sigma_v \simeq 0.42 \left(\frac{M}{\mathrm{M}_\odot}\right)^{0.2} \ \mathrm{km\,s^{-1}} \ ,$$

The final relation involves the cloud average H_2 number density, $\langle n_{H_2} \rangle$, and its size, L:

$$\langle n_{H_2} \rangle \simeq 3400 \left(\frac{L}{1\,\mathrm{pc}} \right)^{-1.1} \ \mathrm{cm}^{-3} \ .$$

Larsons relations: Implications

From the 1st and the 2nd one can infer

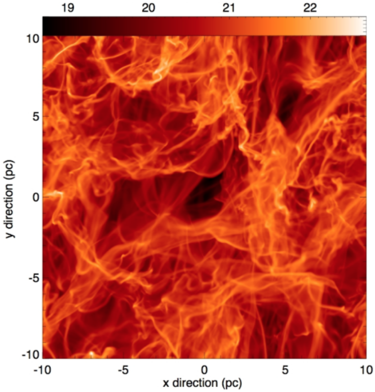
$$\langle\Sigma\rangle \propto ML^{-2} \propto L^{\gamma_L/0.2-2} \ ,$$

GMCs are observed to have roughly constant surface density and the measured values for $\langle\Sigma\rangle$ range from 50 to 400 $\mathrm{M_\odot\ pc^{-2}}$ (see Tan et al. 2013 for a recent review).

The 1st relation implies that (if $L \gtrsim 0.02$ pc) $\sigma_v > c_{\rm th}$, i.e. the gas within cloud is supersonic.

Larsons relations: Implications

Supersonic flows dominate in GMCs and are able to generate a very large density contrast within a cloud up to several order of magnitudes. (Glover et al. 2007)



(see Kritsuk et al. 2013 for a more extended discussion)

Larsons relations: Implications

If clouds are dominated by (supersonic) *turbulence* it is possible to show that the first Larson relation naturally follows.

The kinetic energy for a cloud of size L is proportional to $\sigma_v(L)^2$, where $\sigma_v(L)$ is the velocity dispersion measured on the same scale. If the velocity power spectrum is $P(k) \propto k^{-n_{\rm turb}}$, one has

$$\sigma_v(L)^2 \propto \int_{2\pi/L}^{\infty} \mathrm{d}k\, P(k) \propto L^{n_{\rm turb}-1} \quad .$$

For incompressible, supersonic turbulence, it can be shown that $n_{\rm turb} = 2$, which leads to $\sigma_v \propto L^{0.5}$.

Larsons relations: Implications

The first Larson relation can also be recovered if clouds are virialized. For example, this happens if their virial parameter

$$\alpha_{\rm vir} \equiv \frac{2\mathcal{K}}{|\mathcal{W}|} \simeq 1 \ ,$$

where $\mathcal{K} = 3/2\, M\, \sigma_v^2$ and $\mathcal{W} \simeq -3/5\, G\, M^2/L$ are the cloud kinetic and gravitational energy, respectively.

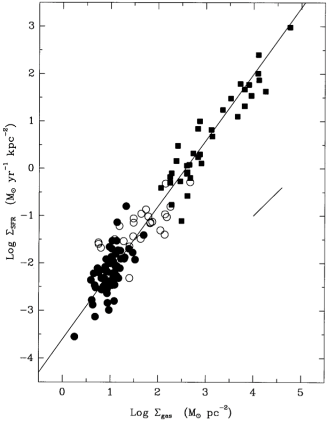
Most observed clouds have $\alpha_{\rm vir}$ in the range $0.5 - 3$ (see Krumholz et al. 2007), thus this implies

$$\sigma_v \propto \sqrt{\pi G \Sigma} L^{1/2} \ ,$$

which is equivalent to the first Larson relation when Σ is constant.

Star Formation: The Kennicutt Relation

$$\left(\frac{\Sigma_{\rm SFR}}{M_\odot\,{\rm yr}^{-1}\,{\rm kpc}^{-2}}\right) = 2.5\times 10^{-4}\left(\frac{\Sigma_{\rm gas}}{M_\odot\,{\rm pc}^{-2}}\right)^{1.5}$$



(Kennicutt et al. 1989, Kennicutt et al. 1998)

Star Formation: The Kennicutt Relation

Usually explained in terms of an underlying *Schmidt relation* $\dot{\rho}_* = \varepsilon_{\rm ff} \rho_{\rm gas}/t_{\rm ff}$, where

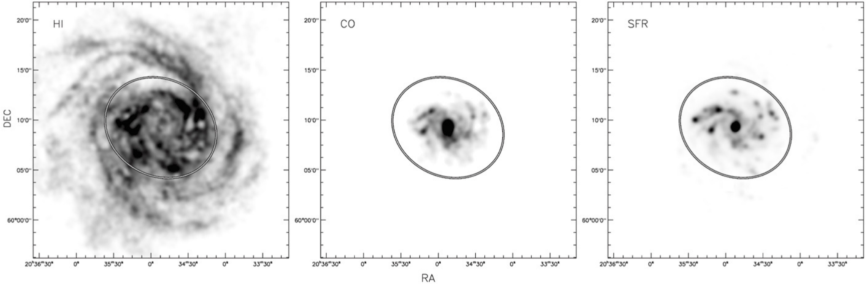
$\rho_{\rm gas}$ is the gas density

$$t_{\rm ff} = \sqrt{\frac{3\pi}{32\,G\,\rho_{\rm gas}}} \simeq 5.1 \times 10^6 \left(\frac{n_{\rm H}}{100\,{\rm cm}^-3}\right)^{-1/2} {\rm yr}$$

$\varepsilon_{\rm ff} \simeq 0.01$ is the star formation efficiency per free-fall time, observed to be nearly constant in GMCs (Kennicutt et al. 2007)

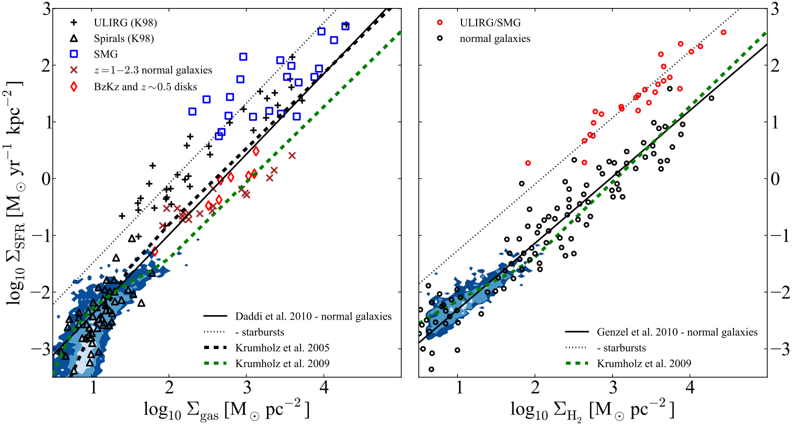
Therefore $\dot{\rho}_* \propto \rho_{\rm gas}^{1.5}$

H_2 and Star Formation



(Bigiel et al. 2008)

H_2 and Star Formation



H_2 and Star Formation

H_2 and star formation correlate when on galactic scales (i.e. at the half-light radius) as well as on local scales ($\sim$ 300 pc, see Kennicutt et al. 2007, Bigiel et al. 2008)

The amplitude
and the slope of the relation does not seem to depend on the scale

All of this has been interpreted as this correlation to be *fundamental, universal*.

Can the observed correlation be recovered from first principles?

If GMCs are dominated by supersonic turbulence one can compute the fraction of the total mass of a cloud unstable to gravitational collapse as

$$f_J = \int_{x_J}^{\infty} x\mathcal{P}(x)\mathrm{d}x$$

where $x_J = n_J/\overline{n}$, with n_J being the Jeans density for a cloud of mass M , and $\mathcal{P}(x)$ being the density probability distribution function inside GMCs (i.e. it can be approximated with a *lognormal distribution*, see Padoan et al. 1997, Kainulanen et al. 2009).

Can the observed correlation be recovered from first principles?

Starting from this equation Krumholz et al. 2005 gave an expression for the star-formation efficiency per free-fall time

$$\varepsilon_{\rm ff} = \frac{\epsilon_{\rm core}}{\phi} f_J \quad ,$$

where $\epsilon_{\rm core}$ is the fraction of the collapsing mass f_J that eventually ends up in stars and $\phi\,(>1)$ takes into account that star formation takes place on time scales larger than the free-fall time. After further simplifications, they showed that

$$\varepsilon_{\rm ff} \approx 0.014 \left(\frac{\alpha_{\rm vir}}{1.3}\right)^{-0.68} \left(\frac{\mathcal{M}}{100}\right)^{-0.32} \quad .$$

Can the observed correlation be recovered from first principles?

If star formation resides in GMCs, which constitute a fraction $f_{\rm GMC}$ of the total gas surface density, $\Sigma_{\rm gas}$, the star-formation rate density for a disk can be expressed as

$$\Sigma_{\rm SFR} = \frac{\varepsilon_{\rm ff} f_{\rm GMC} \Sigma_{\rm gas}}{t_{\rm ff}} \quad .$$

Under the assumptions that clouds are virialized and their surface density is determined by the external pressure in a galaxy it can be shown that (Krumholz et al. 2005)

$$\Sigma_{\rm SFR} \simeq 0.16 \frac{\Sigma_{\rm gas,2}^{1.33}}{1 + 0.025\,\Sigma_{\rm gas,2}^{-2}}$$

where $\Sigma_{\rm gas,2} = \Sigma_{\rm gas}/100\,{\rm M_\odot\,pc^{-2}}$

Why the observed relation shows two different slopes?

i) There are two different X_{CO} conversion factors for normal galaxies and mergers / starbursts

ii) A fundamental law is a volumetric one and the dichotomy is only a projection effect

i) is commonly accepted, because the density in starburst is about 100 times higher than in normal galaxies (see Bolatto et al. 2013 for a recent review). Hence the CO luminosity is higher in those galaxies for a given H_2 mass.

ii) Krumholz, Dekel and McKee 2011 proposed that a local, volumetric star formation relation is consistent with observations. The variations stem from the fact that observed objects have a large range of 3D size scales.

H_2 and SF: correlation or causation?

Molecular hydrogen is only found in the CNM, in well-shielded regions against interstellar UV radiation, where the gas thermal pressure is low and favors gravitational collapse. Therefore, the observed correlation between star formation and molecules might just be a signature of the much deeper relation between temperature and chemical state. However, at sufficiently low gas metallicity, when H_2 formation on dust grains becomes very inefficient, the correlation should break down.

H_2 and SF: correlation or causation?

For instance, if we consider a GMC with mean hydrogen number density $\langle n_{\rm H} \rangle$, average temperature T and metallicity Z and compared three time scale

i) the thermal equilibrium time scale ($t_{\rm cool}$)
ii) the time to reach chemical equilibrium ($t_{\rm H_2}$)
iii) the free-fall time ($t_{\rm ff}$)

If $t_{\rm cool}/t_{\rm ff} \ll 1$ and $t_{\rm H_2}/t_{\rm ff} \ll 1$ the cloud can cool and form molecules rapidly, hence we expect a correlation between H_2 and star formation.

If $t_{\rm cool}/t_{\rm ff} \ll 1$ but $t_{\rm H_2}/t_{\rm ff} \gg 1$, the gravitational collapse will still occur but H_2 formation will proceed very slowly, and no correlation between molecules and stars is expected.

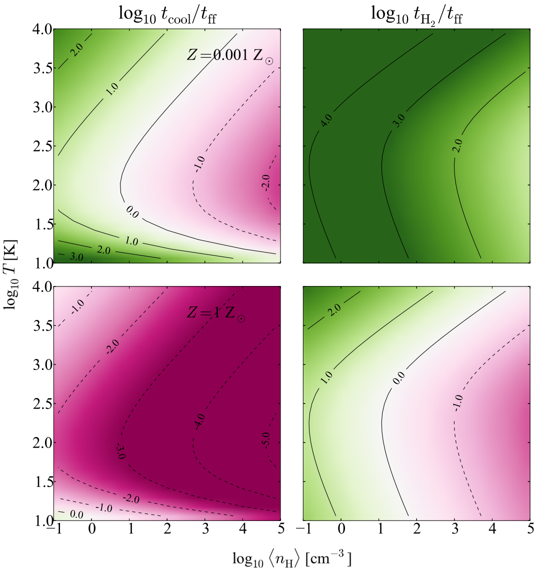
H_2 and SF: correlation or causation?

Following Krumholz et al. 2012 we have

$$\frac{t_{\rm cool}}{t_{\rm ff}} \sim 8.3 \times 10^{-4} \left(\frac{T}{T_{\rm C\,II}}\right) e^{T_{\rm C\,II}/T} \left(\frac{Z}{Z_\odot}\right)^{-1} \left(\frac{C_\rho}{10}\right)^{-1} \left(\frac{\langle n_{\rm H}\rangle}{1\,{\rm cm}^{-3}}\right)^{-1/2},$$

$$\frac{t_{\rm H_2}}{t_{\rm ff}} \sim 2.4\, S_{\rm H}(T)^{-1} \left(\frac{T}{100\,{\rm K}}\right)^{-1/2} \left(\frac{Z}{Z_\odot}\right)^{-1} \left(\frac{C_\rho}{10}\right)^{-1} \left(\frac{\langle n_{\rm H}\rangle}{1\,{\rm cm}^{-3}}\right)^{-1/2}.$$

H_2 and SF: correlation or causation?



H_2 and SF: correlation or causation?

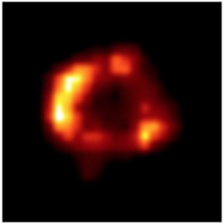
Glover et al. 2012 approached a similar problem with hydrodynamical simulations of resolved molecular clouds, that included a complex chemical network and molecular cooling functions.

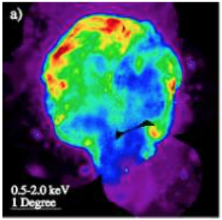
They showed that metallicity has very little effect on the star formation in the cloud, with metal-poor clouds experiencing a delay in the onset of star formation by no more than a cloud free-fall time. Furthermore, the presence of H_2 hardly influences the evolution of the cloud and it contributes as little as 1 per cent to the total cooling rate. On the other hand, if one disables the effect of dust shielding, the gas in the cloud cannot cool below 100 K and this results in a strong suppression of star formation.

H_2 and SF: correlation or causation?

A simple way to verify this prediction is to look at regions that show a high star-formation rate, but a negligible molecular fraction. However, resolved studies of star-forming regions, are only possible in the Milky Way and for a few nearby galaxies, where the conditions are favourable enough for molecular hydrogen to be in chemical equilibrium. In galaxies with metallicities $\lesssim 0.1\ Z_\odot$, CO breaks down as an H_2 tracer.

Supernova Remnants





i) **Free-expansion**: *Chandra* X-ray image of 1987A

ii) **Sedov-Taylor phase**: HST view of the Crab Nebula

iii) **Snow-plow phase**: Rosat X-Ray image of Cygnus Loop

Evolution of a Supernova Remnant

During a SuperNova (SN) explosion a typical mass of 1 $M_\odot$ is ejected into the ISM, with a typical velocity of 10^4 km s^{-1}. This corresponds to a total kinetic energy of

$$E_{\rm kin} = \frac{1}{2} M_{\rm ejecta} V_{\rm ejecta}^2 \simeq 10^{51} {\rm erg} \quad .$$

Evolution of a Supernova Remnant: Free expansion

The first phase a **free-expansion** that roughly ends when the mass swept up from the expanding shell is equal to the mass of the ejecta

$$\frac{4}{3}\,\pi\,r^3\,\rho_0\,M_{\rm ejecta} \quad ,$$

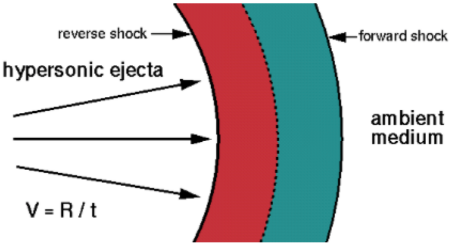
where we have assumed that the ISM is uniform ($\rho_0 = \mathrm{const.}$)

The free-expansion phase ends when

$$r \simeq 2\,\mathrm{pc}\,(\mu\,n_0)^{-1/3}$$

$$t_{\rm f.e.} \simeq \frac{2\,\mathrm{pc}}{10^4\,\mathrm{km\,s^{-1}}} \sim 200\,\mathrm{yr}$$

Evolution of a Supernova Remnant



Now the shock front and a reverse shock have formed and the ejecta slow down.

Evolution of a Supernova Remnant: Adiabatic expansion

How thick is the shell?

Because we are in the present of an adiabatic shock, we have

$$\rho_1 = 4\,\rho_0$$
$$P_1 = \frac{3}{4}\,\rho_0\,v_0^2$$

Therefore, on the assumption that all the mass is swept up into the shell, we have

$$\frac{4\pi}{3}\,\rho_0\,r_s^3 = 4\pi\,\rho_1\,r_s\,D \quad ,$$

we have $D/r_s = 1/12$ which means that the shell is thin.

How do the velocity, energy, and temperature of the shell change with time? **Sedov-Taylor (ST) solution**

Evolution of a Supernova Remnant: Adiabatic expansion

The temperature of the medium immediately surrounding the shock can be estimated to be (for an adiabatic shock with $\gamma = 5/3$)

$$\begin{aligned} T_1 &= \frac{\mu\, m_{\rm H}}{k_B} \frac{P_1}{\rho_1} \\ &= \frac{\mu\, m_{\rm H}}{k_B} \, \frac{3}{4} \rho_0 \, v_0^2 \frac{1}{4\rho_0} \\ &= \frac{3}{16} \frac{\mu\, m_{\rm H}}{k_B} \, v_0^2 \simeq 2 \times 10^9 \, {\rm K} \ , \end{aligned}$$

The cooling time can be estimated to be

$$t_{\rm cool} = \frac{\frac{3}{2} n\, k_B\, T_1}{n_e\, n_i\, \Lambda(T)} \simeq 3 \times 10^8 \, {\rm yr}$$

This means that the expansion proceeds **adiabatically** (i.e. radiative losses are not important)

Evolution of a Supernova Remnant: Sedov-Taylor solution

The equation of motion for an expanding shell is

$$\frac{\mathrm{d}}{\mathrm{d}t}\left(\frac{4\pi}{3}\,\rho_0\,r_s^3\,v_s\right) = 4\pi\,r_s^2\,P \quad ,$$

where

$$P = (\gamma - 1)\frac{E}{V} = \frac{E_{\mathrm{SN}}}{2\pi\,r_s^3} \quad .$$

So we have

$$\frac{\mathrm{d}}{\mathrm{d}t}\left(\frac{1}{3}\,r_s^3\,\rho\,\dot{r}_s\right) = \frac{E_{\mathrm{SN}}}{2\pi\,r_s^3} \quad ,$$

and we can find a power-law solution $r_s = A\,t^{\alpha}$.

Evolution of a Supernova Remnant: Sedov-Taylor solution

One can show that $\alpha = 2/5$ and

$$r_s(t) = \left(\frac{25}{4\pi} \frac{E_{\rm SN}}{\mu\, n_0\, m_{\rm H}} \right)^{1/5} t^{2/5}$$

$$\simeq 0.3 \left(\frac{E_{\rm SN}}{10^{51}\,{\rm erg}} \right)^{1/5} (\mu\, n_0)^{-1/5} \left(\frac{t}{\rm yr} \right)^{2/5} {\rm pc} \quad .$$

Therefore, we have

$$v_s(t) = \frac{2}{5}\left(\frac{25\,E_{\rm SN}}{4\pi\,\rho}\right)^{1/5} \simeq 1.3\times 10^5\ \left(\frac{t}{\rm yr}\right)^{-3/5}\ {\rm yr}$$

$$T_s(t) = \frac{3}{16}\frac{\mu\, m_{\rm H}}{k_B}\, v_s^2$$

$$\simeq 3\times 10^{11}\,\mu^{3/5}\left(\frac{E_{\rm SN}}{10^{51}\,{\rm erg}}\right)^{2/5}\ (n_0)^{-2/5}\ \left(\frac{t}{\rm yr}\right)^{-6/5}\ {\rm pc}$$

When does the adiabatic phase end?
When the cooling time is comparable to the age of SN remnant.

This happens at $t_{\mathrm{S-T}} \simeq 3 \times 10^4$ yr, that corresponds to a temperature of the shell $T_s \simeq 10^6$ K.

Some of the energy was still radiated away. But how much?

$$E_{\mathrm{irr}} \sim n^2 \, \lambda(T) \, v_s \, t_{\mathrm{S-T}} \simeq 10^{49} \, \mathrm{erg} \quad ,$$

which is only 6% of the SN energy budget.

Evolution of a Supernova Remnant: Radiative phase

At $T \sim 10^6$ K atoms start to capture free electrons and cooling starts to dominate. As a consequence, the thermal pressure in the post-shock region decreases and the expansion slows down.

$$\frac{\mathrm{d}}{\mathrm{d}t}\left(\frac{4\pi}{3}\,\rho_0\,r_s^3\,v_s\right) = 4\pi\,r_s^2\,P_b \;\;,$$

Because radiative losses are more significant within the shell, we can assume that inside the bubble the expansion is still adiabatic, so we can write

$P_b = P_{b,A}\,\dfrac{\rho_b^\gamma}{\rho_b,\,A^\gamma}$, where A refers to the physical quantities computed at the end of the adiabatic phase

Evolution of a Supernova Remnant: Radiative phase

We can assume that $M_b/M_{b,A} \simeq 1$, hence

$$P_b = P_{b,A} \left(\frac{r_{s,A}}{r_s}\right)^{3\gamma}.$$

After some algebra one can show that

$$r_s(t) = 1.3 \left(\frac{E_{\mathrm{SN}}}{10^{51}\,\mathrm{erg}}\right)^{1/7} (\mu\, n_0)^{-1/7} \left(\frac{t}{\mathrm{yr}}\right)^{2/7} \mathrm{pc}$$

$$v_s(t) = 4 \times 10^5 \left(\frac{E_{\mathrm{SN}}}{10^{51}\,\mathrm{erg}}\right)^{1/7} (\mu\, n_0)^{-1/7} \left(\frac{t}{\mathrm{yr}}\right)^{-5/7} \mathrm{km\,s^{-1}}$$

Evolution of a Supernova Remnant: Snow plow phase

Eventually the pressure is nearly zero and the SNR enter the so-called *snow plow* phase.

$$\frac{\mathrm{d}}{\mathrm{d}t}\,(M\,V) = 0$$

This means

$$r_s^3\,v_s = r_{s,\mathrm{S-T}}^3\,v_{s,\mathrm{S-T}}$$

so

$$r_s(t) = \left(4\,r_{s,\mathrm{S-T}}^3\,v_{s,\mathrm{S-T}}\right)^{1/4}\,t^{1/4} \quad ,$$

Eventually the shell breaks up into individual clumps, probably due to the Rayleigh-Taylor instability

Evolution of a Supernova Remnant: Snow plow phase

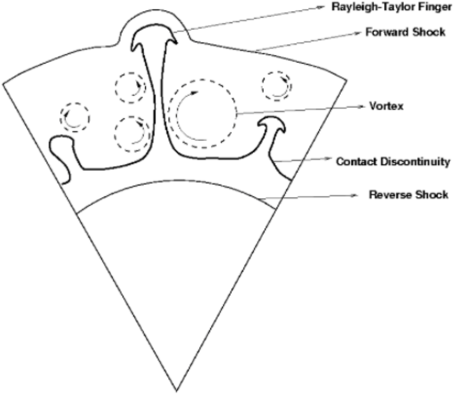
The expansion ends when $v_s \sim 10$ km s^{-1}, i.e. there is no shock

This happens at $t \sim 2 \times 10^6$ yr and the size of shock is nearly $r_s \simeq 80$ pc.

Evolution of a Supernova Remnant: Other considerations

If the medium is not homogeneous?

The shell can be slowed down when it meets dense cloud in the ISM



Stellar Winds

Massive O-B stars have very strong winds, with a mass-loss rate $\dot{M} \sim 10^{-6}$ M$_\odot$ yr^{-1} and a typical wind velocity $v_w \sim 2 \times 10^3$ km s^{-1}. This correspond to a wind mechanical luminosity

$$L_w = \frac{1}{2}\dot{M}v_w^2 \simeq 10^{36}\,\mathrm{erg\,s^{-1}}$$

The wind creates a shock that propagates in the medium.
What are the phases of the stellar wind?

From the conservation of momentum we have

$$\frac{\mathrm{d}}{\mathrm{d}t}\left[\frac{4}{3}\pi r^2 \rho_0 \dot{r}\right] = 4\pi r^2 P_b$$

and the conservation of energy within the bubble

$$\frac{\mathrm{d}U_b}{\mathrm{d}t} = L_w - P_b \frac{\mathrm{d}V}{\mathrm{d}t}$$

One can show that

$$r \simeq 7 \times 10^{-3} \left(\frac{L_w}{10^{36}\,\mathrm{erg\,s^{-1}}}\right)^{1/5} (\mu\, n_0)^{-1/5} \left(\frac{t}{\mathrm{yr}}\right)^{3/5} \;\mathrm{pc}$$

Stellar Winds and Superbubbles

Stellar wind around O-B stars generate bubble of the size of nearly 20-50 pc.

75% of O stars are in associations, known as *O-B association*, which are groups of ~ 100 stars in a very small region (1-10 pc).

The initial phases are similar to the stellar wind, but with an increases wind luminosity $L_w = 100 \times 10^{36}$ erg s^{-1}. After $\sim 10^6$ yr the first SNe start to explode and the energy budget of the superbubble is
$E_{\mathrm{SN}}\, N_{\mathrm{SN}}/t_{\mathrm{OB}} = 10^{51}\,\mathrm{erg} \cdot 10^2/(5 \times 10^7\,\mathrm{yr}) \simeq 10^{38}$ erg s^{1-}.

The radius of a superbubble at $t = 10^6$ yr is ~ 70 pc, which is comparable to the scale height of the disc $\longrightarrow$ *blowout*, *galactic fountain*

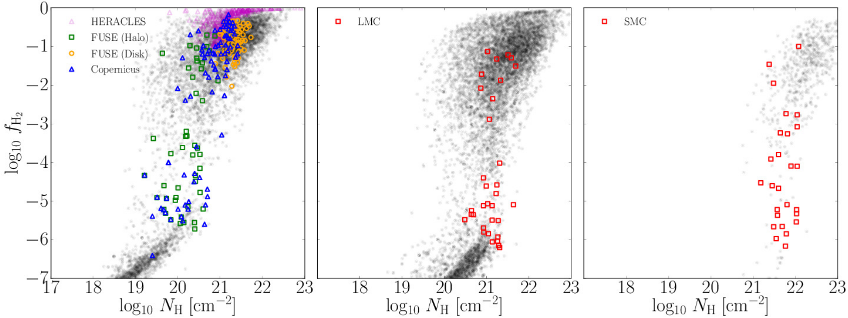
Cross promotion: MoMoGal project

MoMoGal (MOdeling MOlecular hydrogen in GALaxies) project

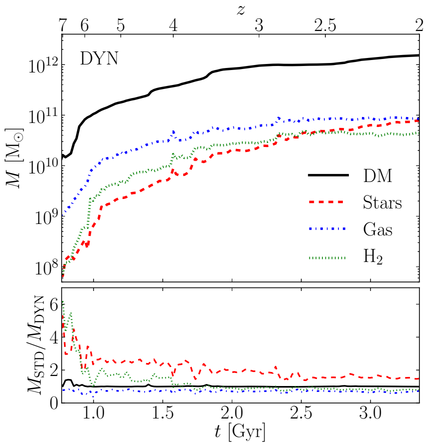
M. Tomassetti et al. 2014, MNRAS, 445 L

M. Tomassetti et al. 2015, MNRAS, 446, 3330T

MoMoGal project



MoMoGal project



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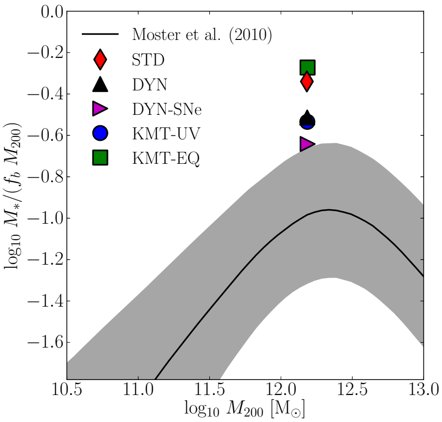
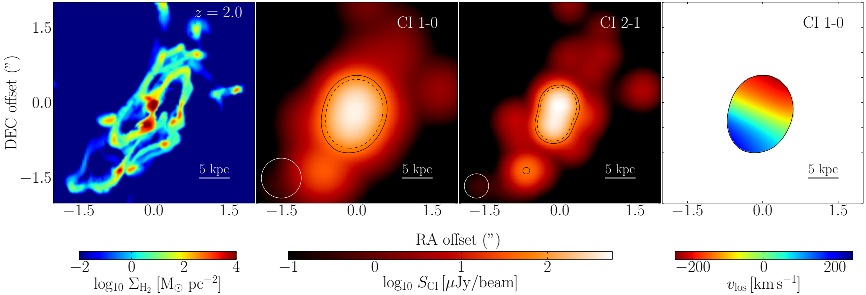


MoMoGal project



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Mark Krumholz's lectures:
https://sites.google.com/a/ucsc.edu/krumholz/teaching-and-courses/ast-230-s-14