Formation and evolution of star-forming molecular clouds

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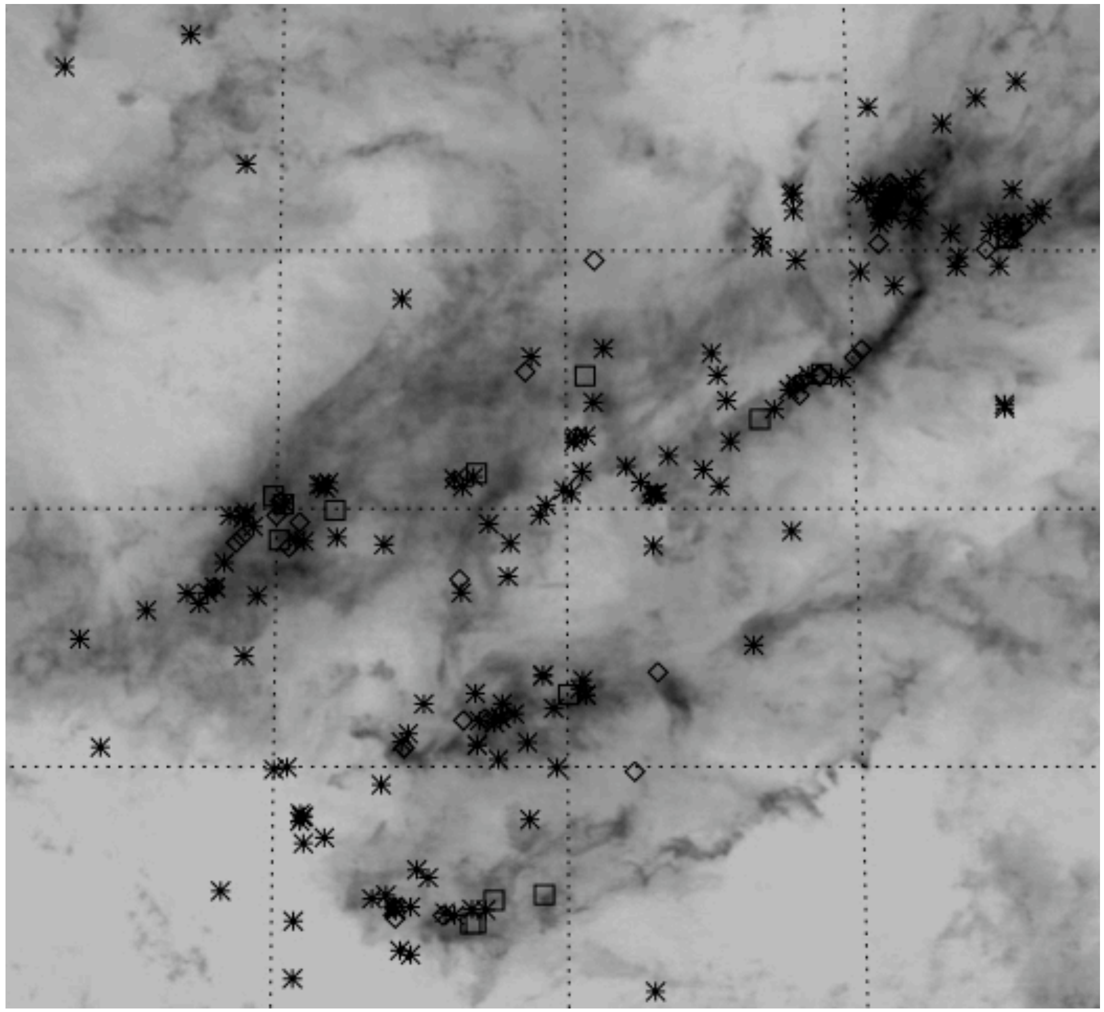
ISIMA 2011

Michigan Star and Planet Formation

molecular clouds:
highly structured ⇒ star formation



Taurus ^{13}CO: Goldsmith et al. 2008

What is a “molecular cloud”?

- Mostly H_2 … shielded from UV heating, T ~ 10 K
- To make H_2;

  – formation; no dipole moment – gas phase formation too slow; ⇒ formation on dust grains; $t(H_2) \sim 10^9$ yr / n(H)

  ⇒ forms in regions with n(H) >~ 10^2 cm^{-3}

  – shielding of dissociating FUV photons by dust and by its own electronic transitions (“self-shielding”)

  ⇒ need column densities N(H) >~ 10^{21} cm^{-3}, $A_V > 0.3$

What is a "molecular cloud"?

• We don't "see" H_2 directly (typically)

• rotational transitions are at relatively high energy ⇒ require higher T to excite (>~ 100 K)

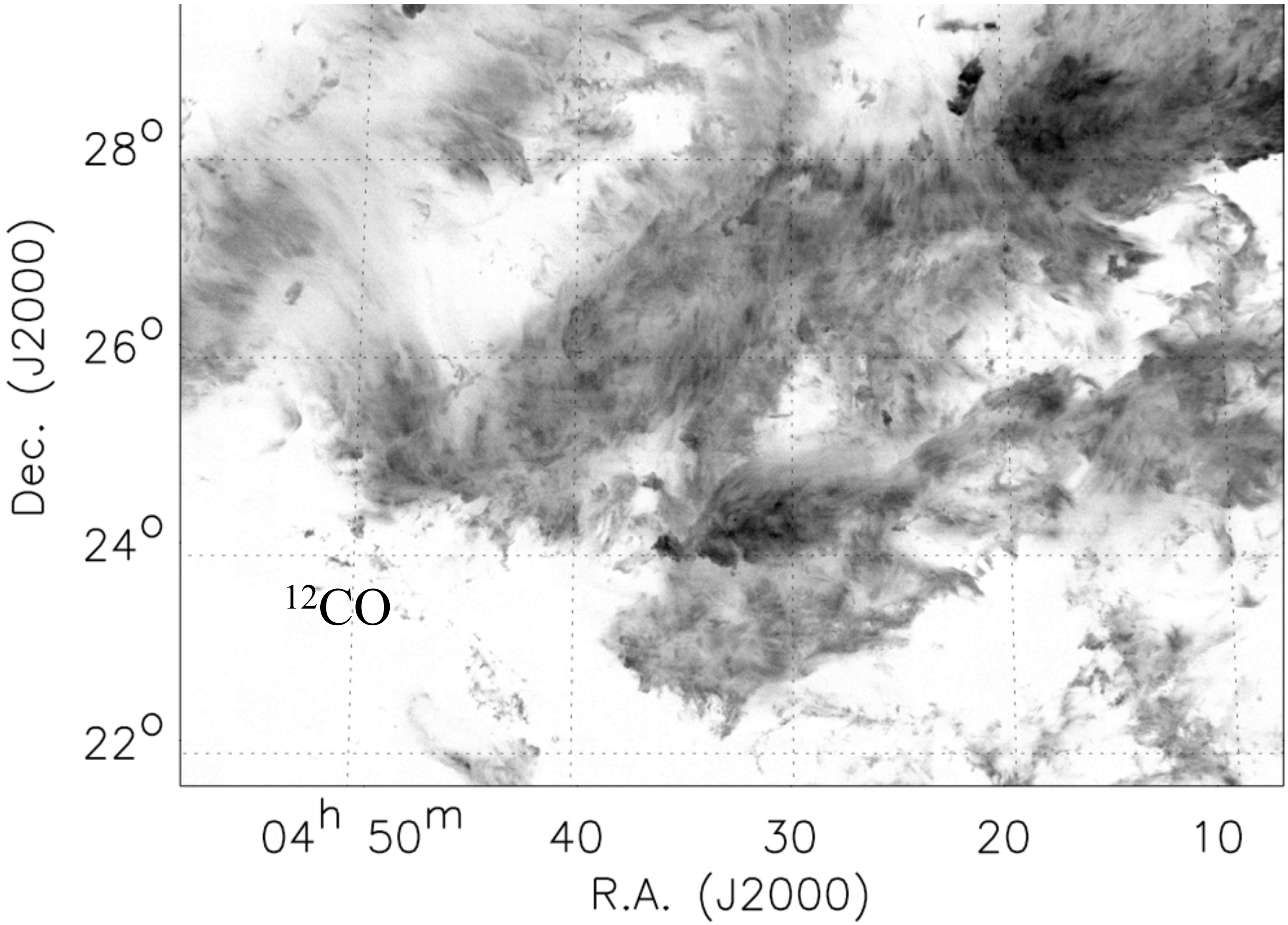
•angular momentum = J ħ; $E = J(J+1)\,\hbar^2/2\,I$

⇒ want heavier molecule with larger moment of inertia I

⇒ CO! (lines at 2.6, 1.3 mm; E/k ~ 5.5 K, 16.5 K)

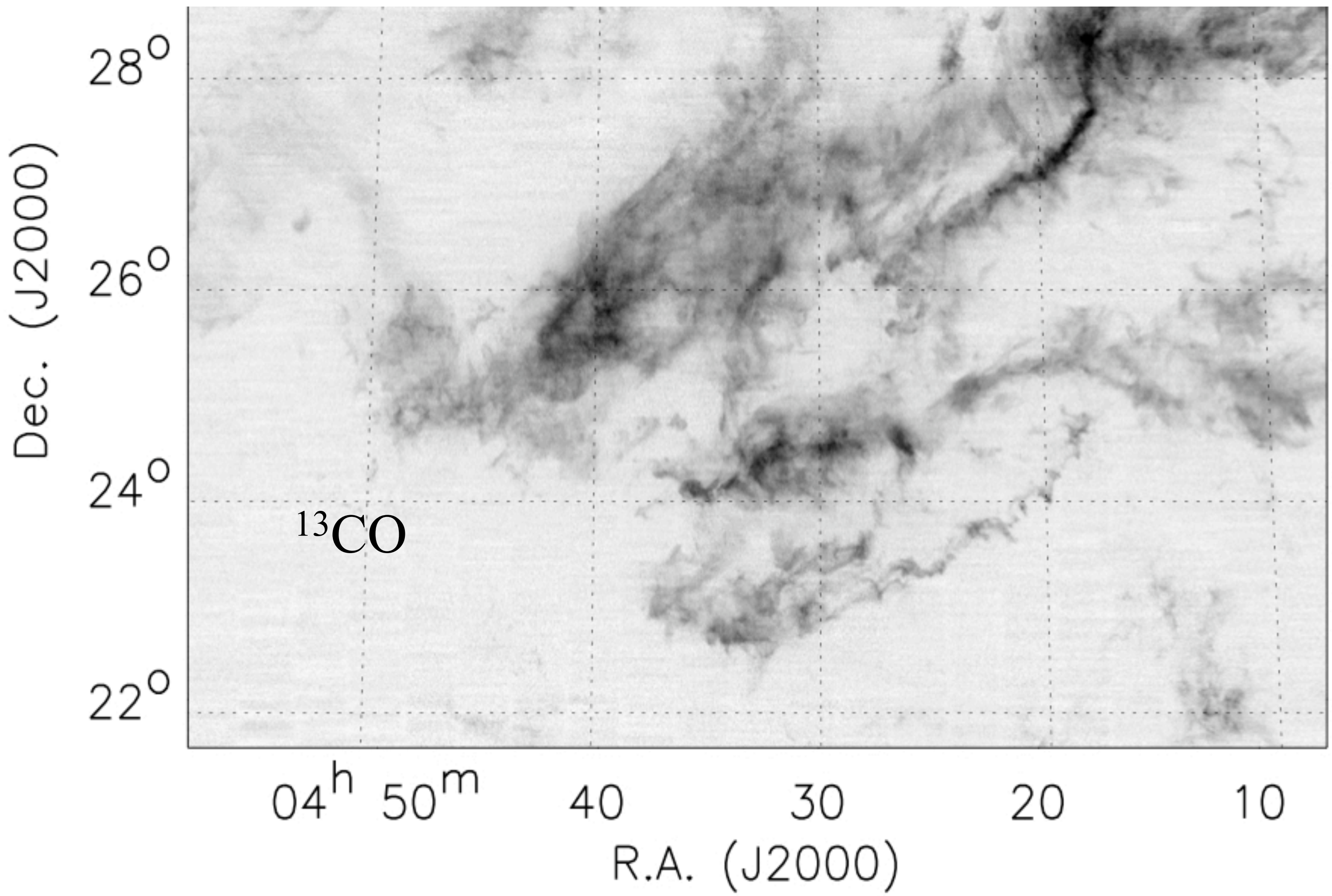
⇒ need density AND shielding; $N_H \sim 3 \times 10^{20}$ cm^{-2} ; A_V >~ 0.5 – 1 (CO dissociates via FUV electronic transitions like H_2)

^{12}CO is usually *optically thick*

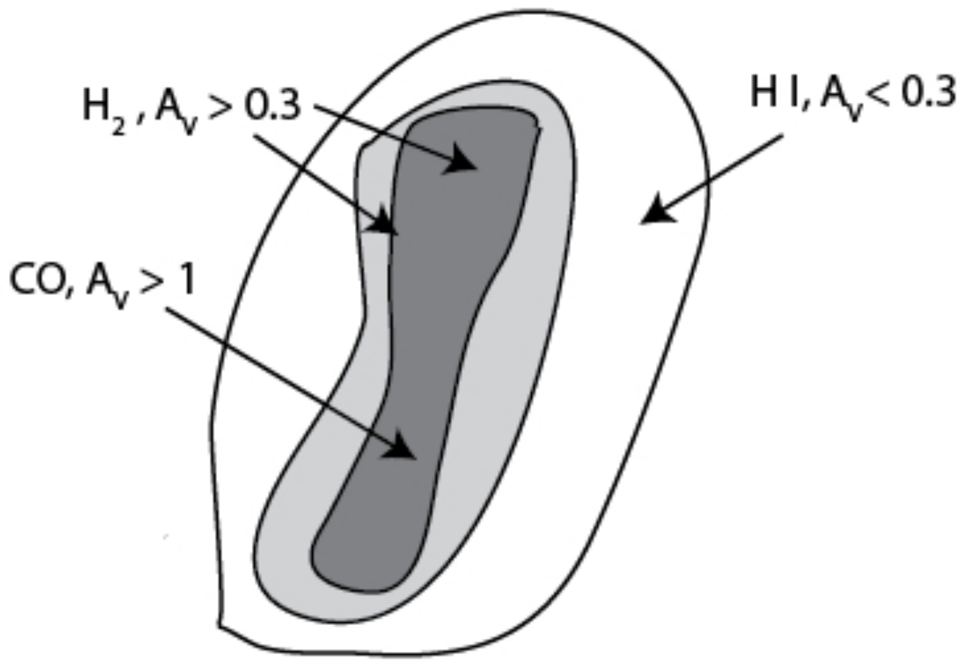


Taurus: Goldsmith et al. 2008

Use isotope ($^{13}CO/^{12}CO$ ratio ~ 1/65) to detect optically-thin regions (and thus measure mass more directly



A typical molecular cloud



Estimating masses not trivial.

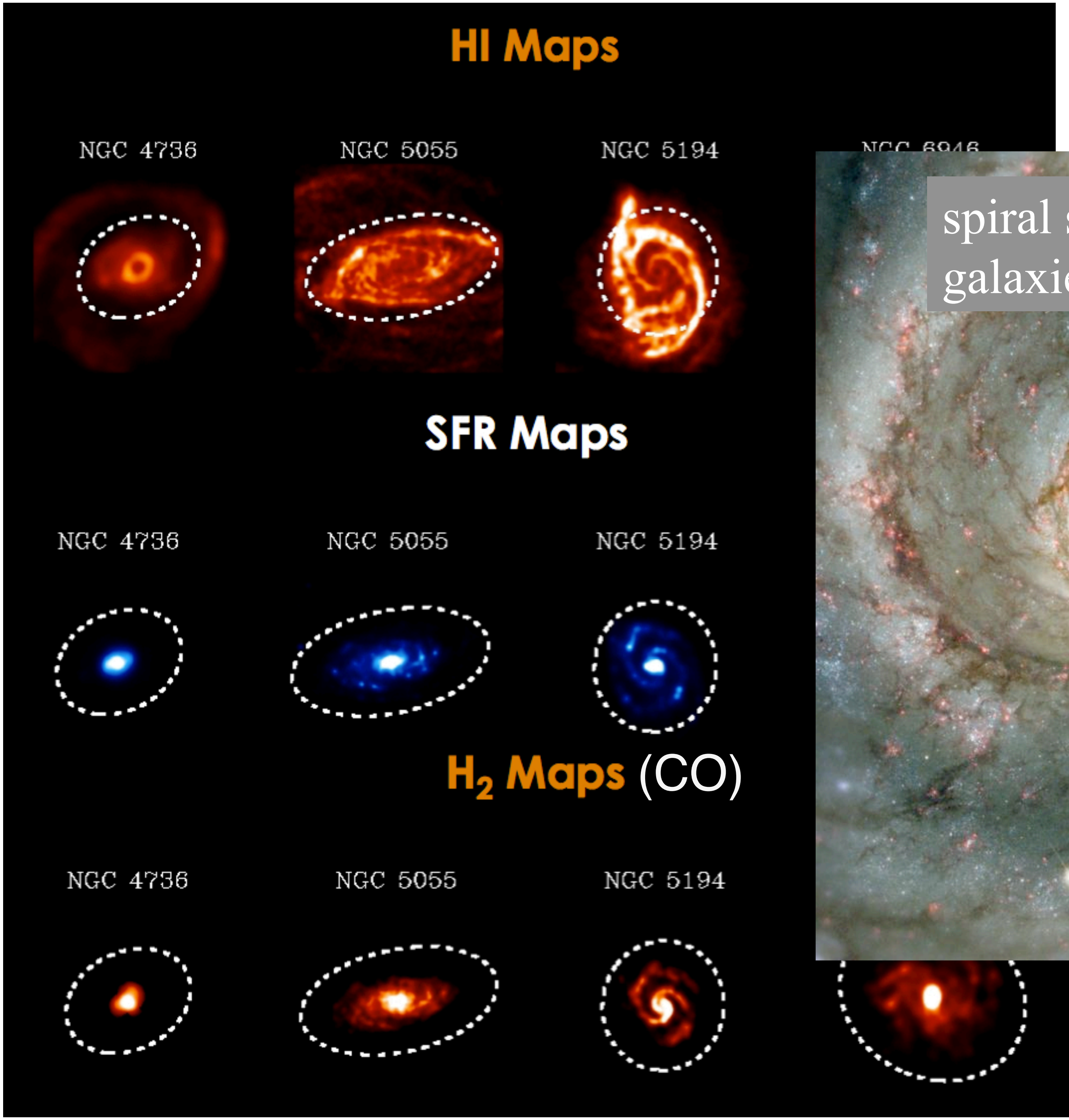
- some H_2 outside CO region
- ^{12}CO 1-0 optically thick;
- ^{13}CO 1-0 more thin, but need isotope ratio to get... excitation temperatures
- bottom line; likely errors of factor of 2 in M(cloud)

Where do molecular clouds come from?

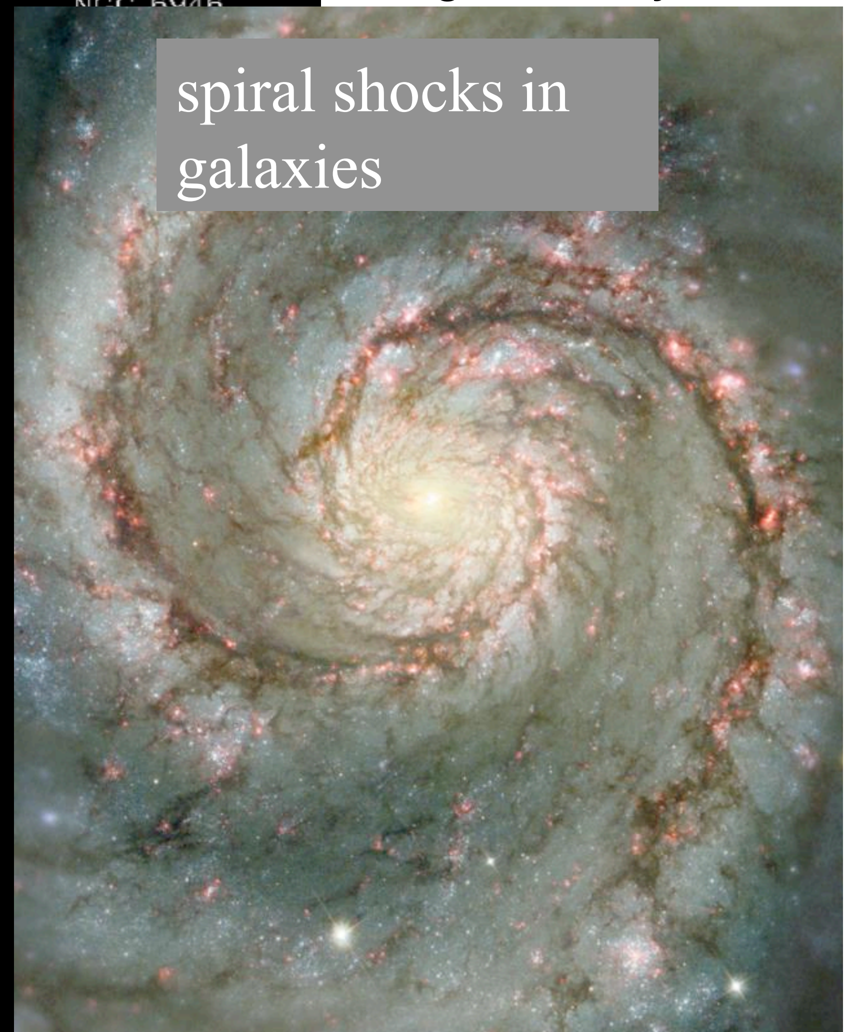
• Accumulation/compression of ISM gas into dense regions with sufficient column density

• mechanisms for sweep-up, compression:

- spiral density waves
- stellar energy input (winds, supernovae, radiation pressure, photoionization, etc.)

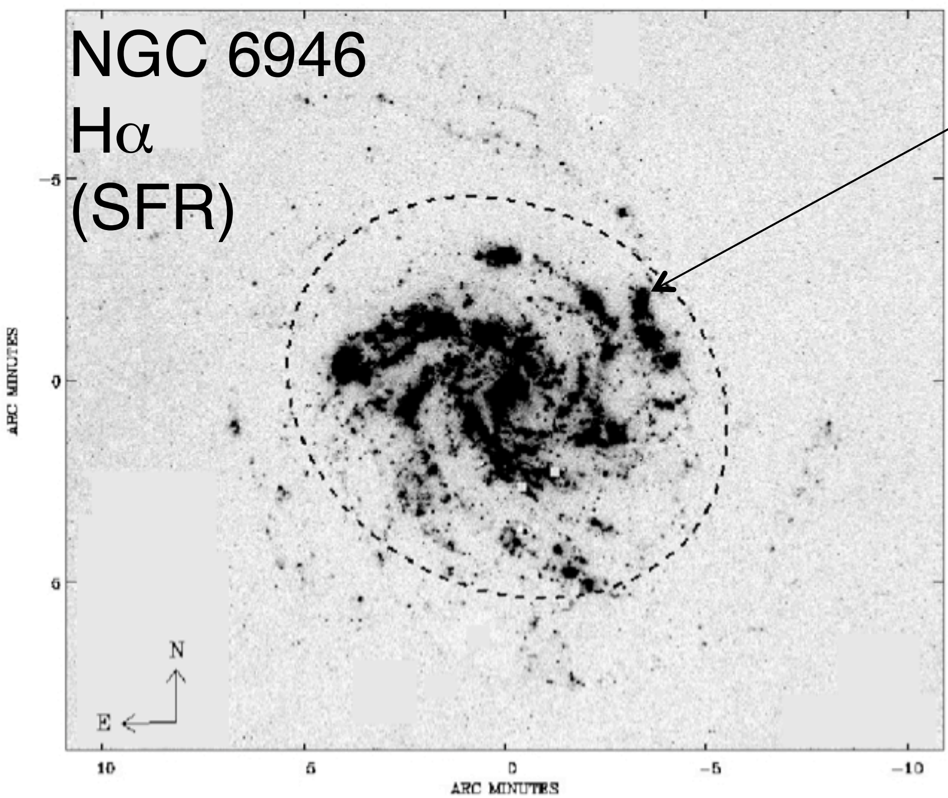


Bigiel, Leroy, et



spiral shocks in galaxies

Being “molecular” just means denser, colder gas;
Cloud/star formation where gas is ***compressed by shocks***

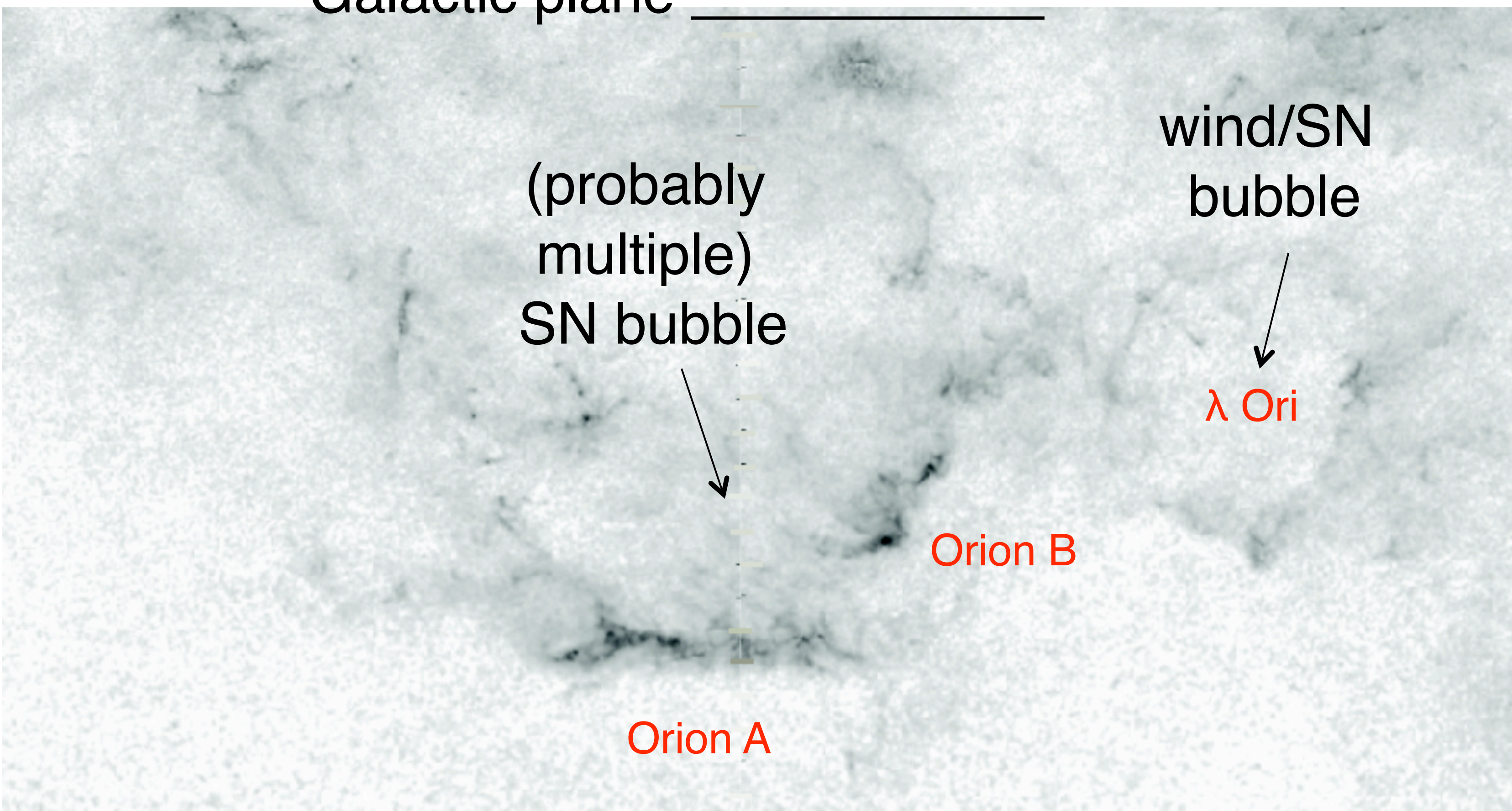


- Inner disk: spiral arms(gravitational instability)

- Outer disk: spiral shock driven by gravitational instability in inner disk.
⇒*compression leads to star formation*

Ferguson et al. 1998

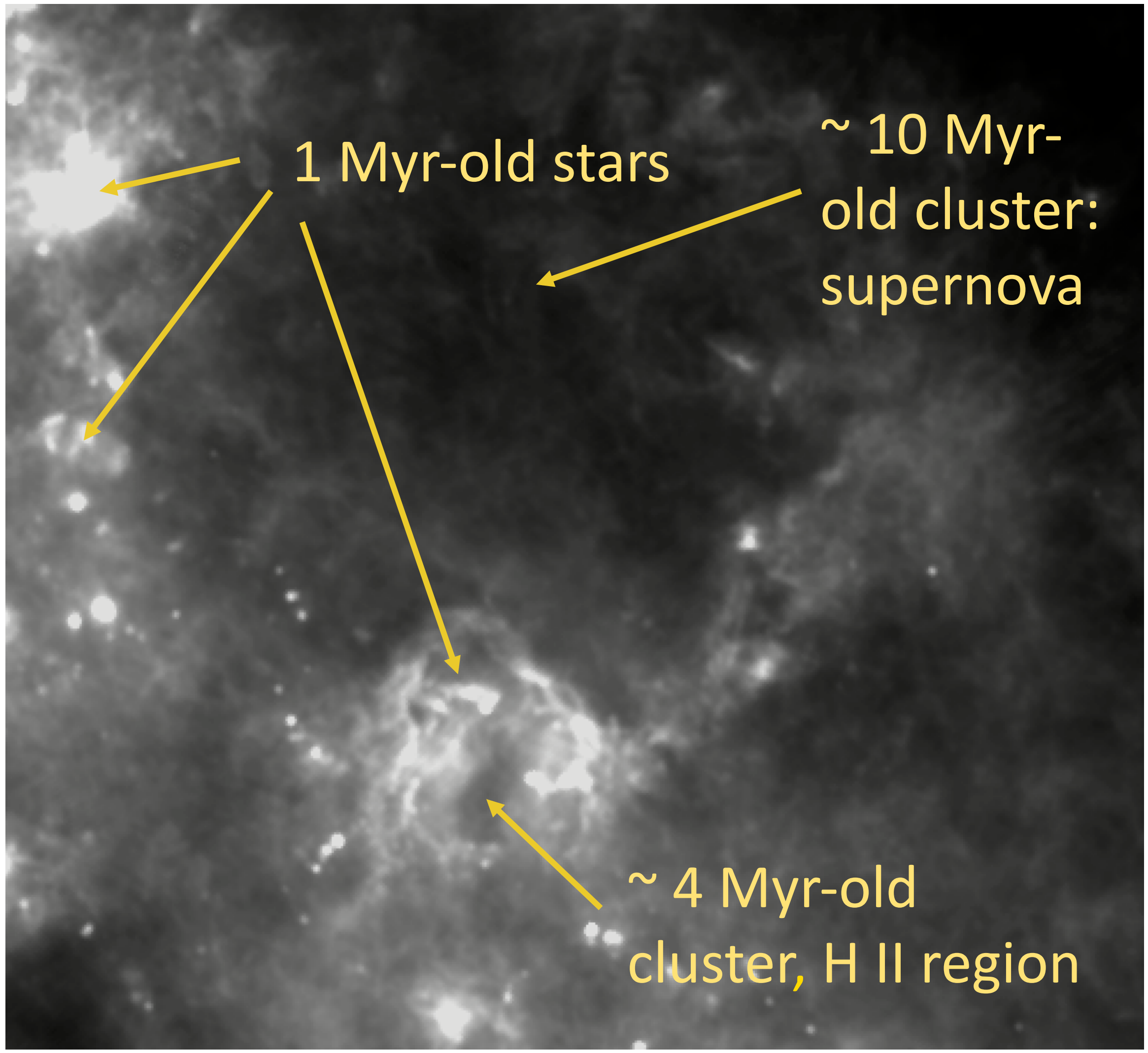
Stellar energy input also forms molecular clouds



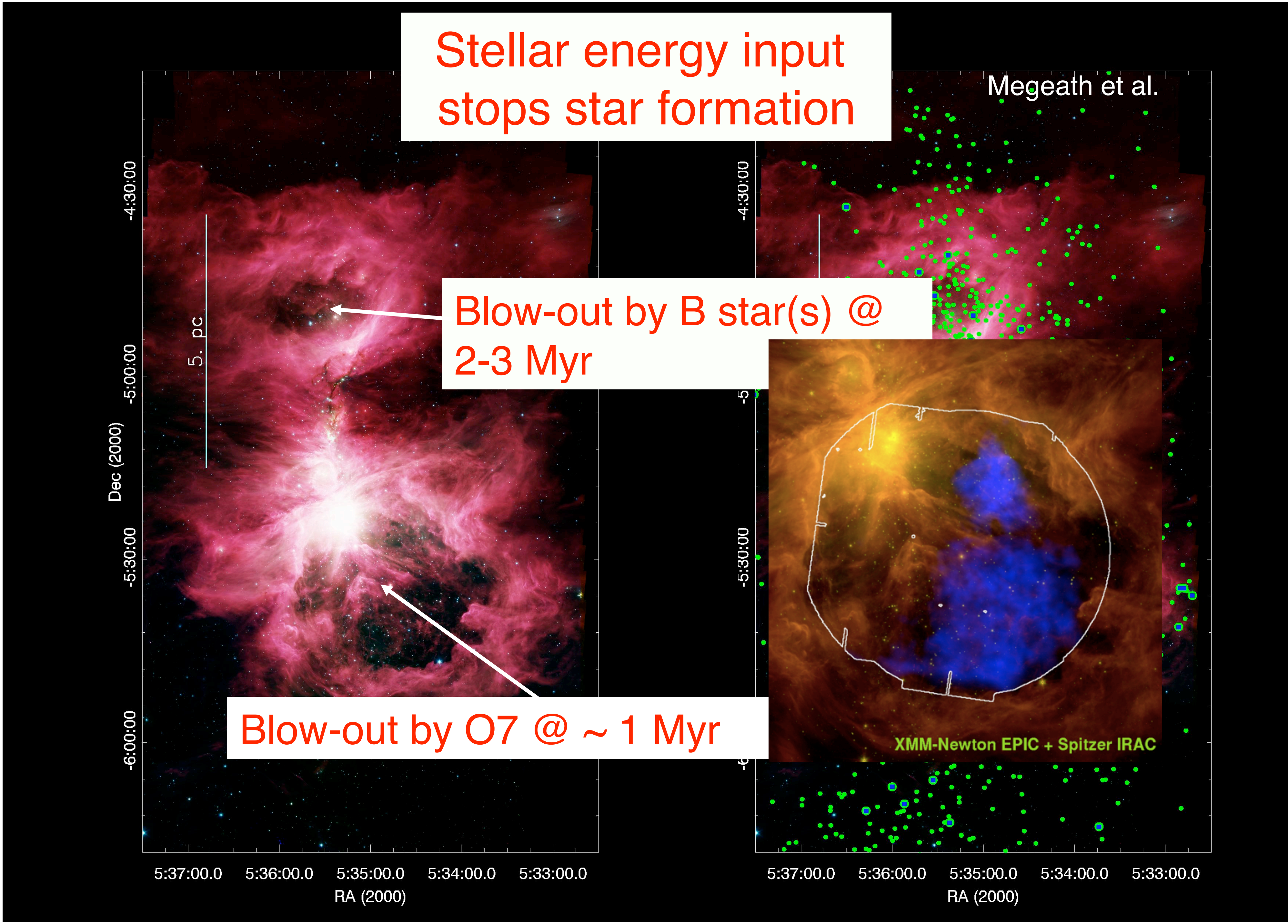
Froebrich & Rowles 2010, A_V map

Cep OB2: supernova, H II region-driven bubbles

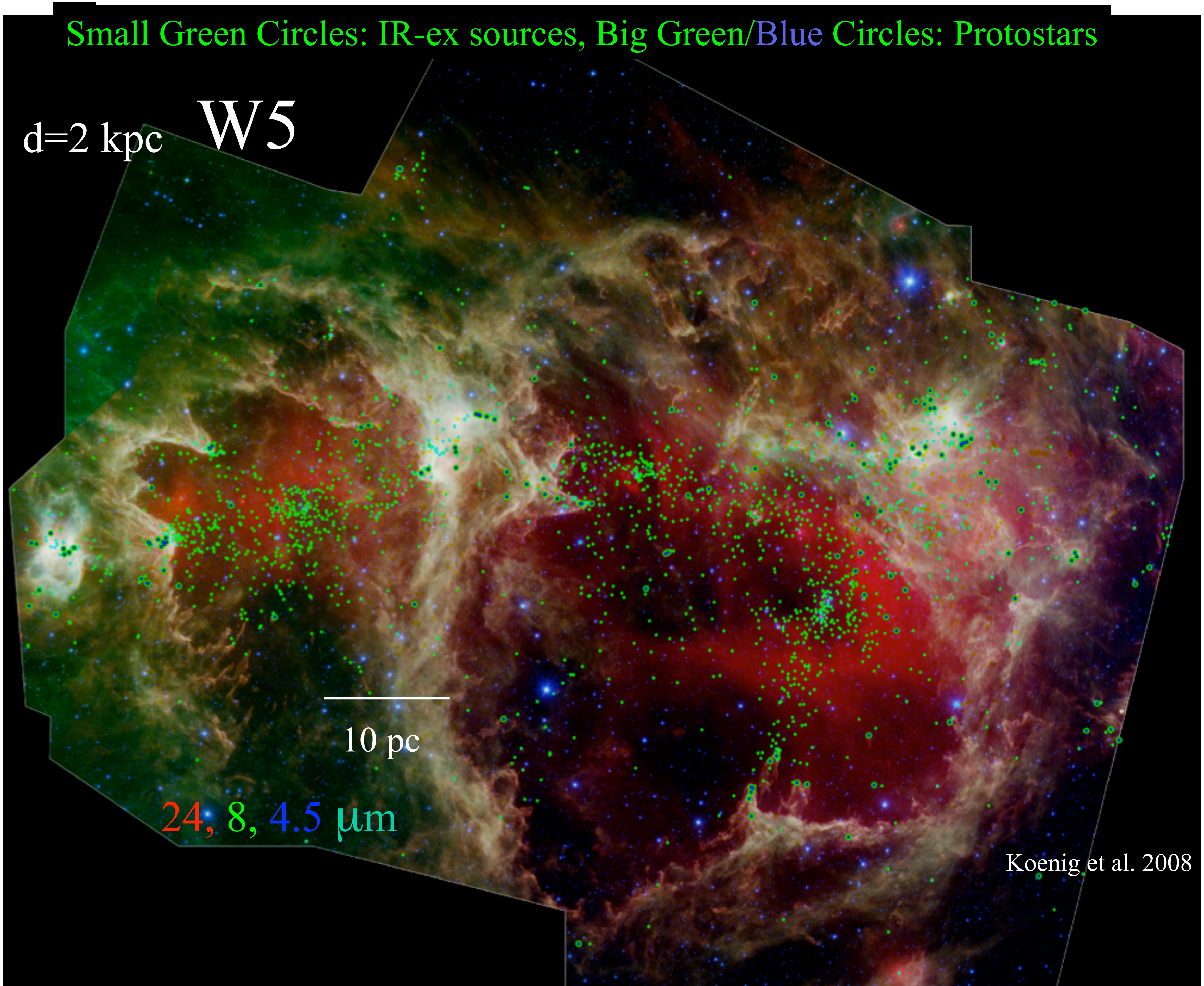
50 pc



100 μm IRAS
dust emission


Stellar energy input
stops star formation
Megeath et al.
Blow-out by B star(s) @
2-3 Myr
Blow-out by O7 @ ~ 1 Myr
XMM-Newton EPIC + Spitzer IRAC
5. pc
Dec (2000)
-4:30:00
-5:00:00
-5:30:00
-6:00:00
5:37:00.0
5:36:00.0
5:35:00.0
5:34:00.0
5:33:00.0
RA (2000)

stellar energy input starts, stops star formation



What are the dynamical states of (dense gas in) molecular clouds?

$$\rho \frac{\partial \mathbf{v}}{\partial t} + \rho \mathbf{v} \cdot \nabla \mathbf{v} = -\nabla \mathbf{P} - \rho \nabla \Phi + \mathbf{F}$$

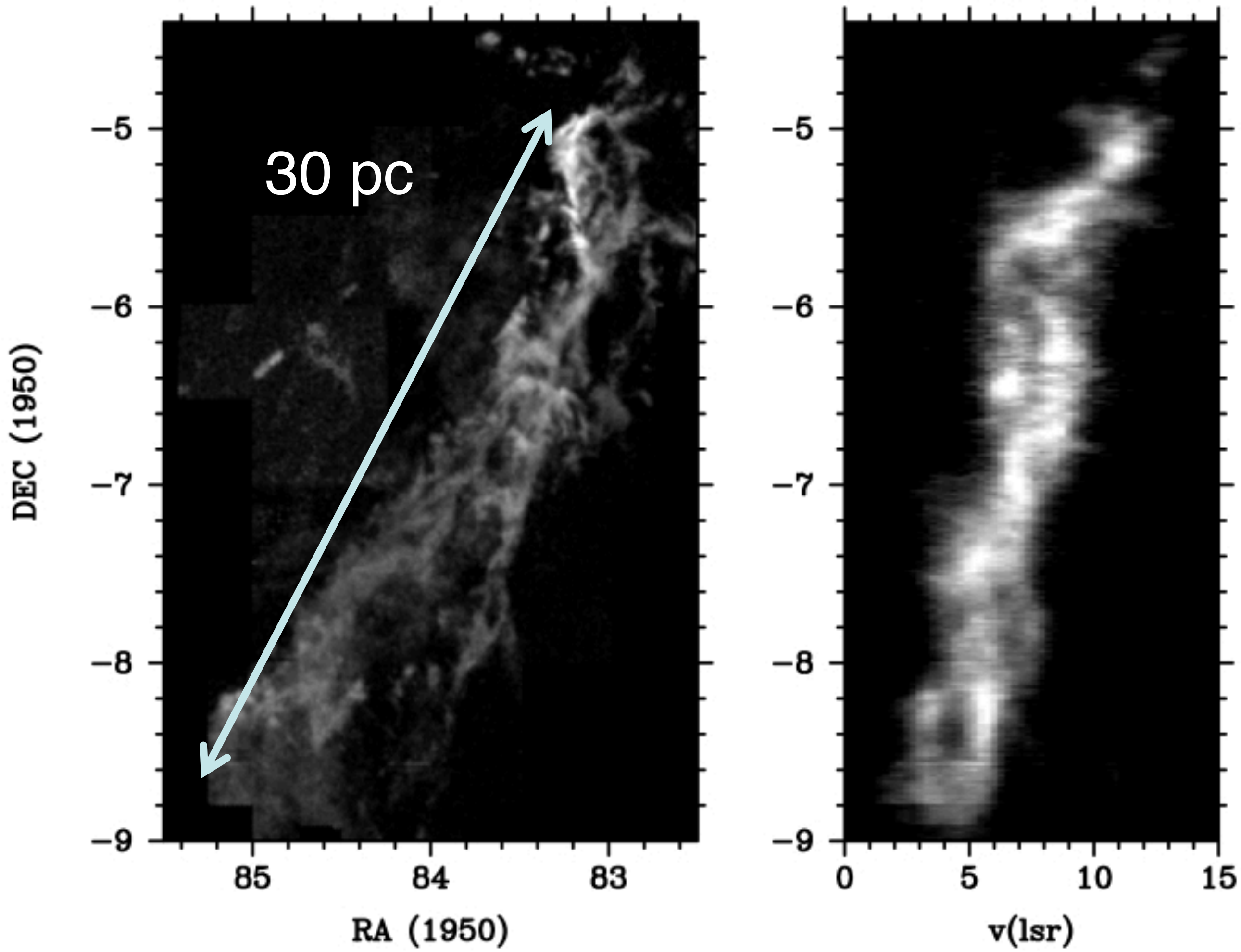
order of magnitude: $v^2 \sim GM/r$ ($\gg c_s^2$)

- take cloud radius ~ 30 pc/2, M ~ 10^5 M(sun) (Orion A)

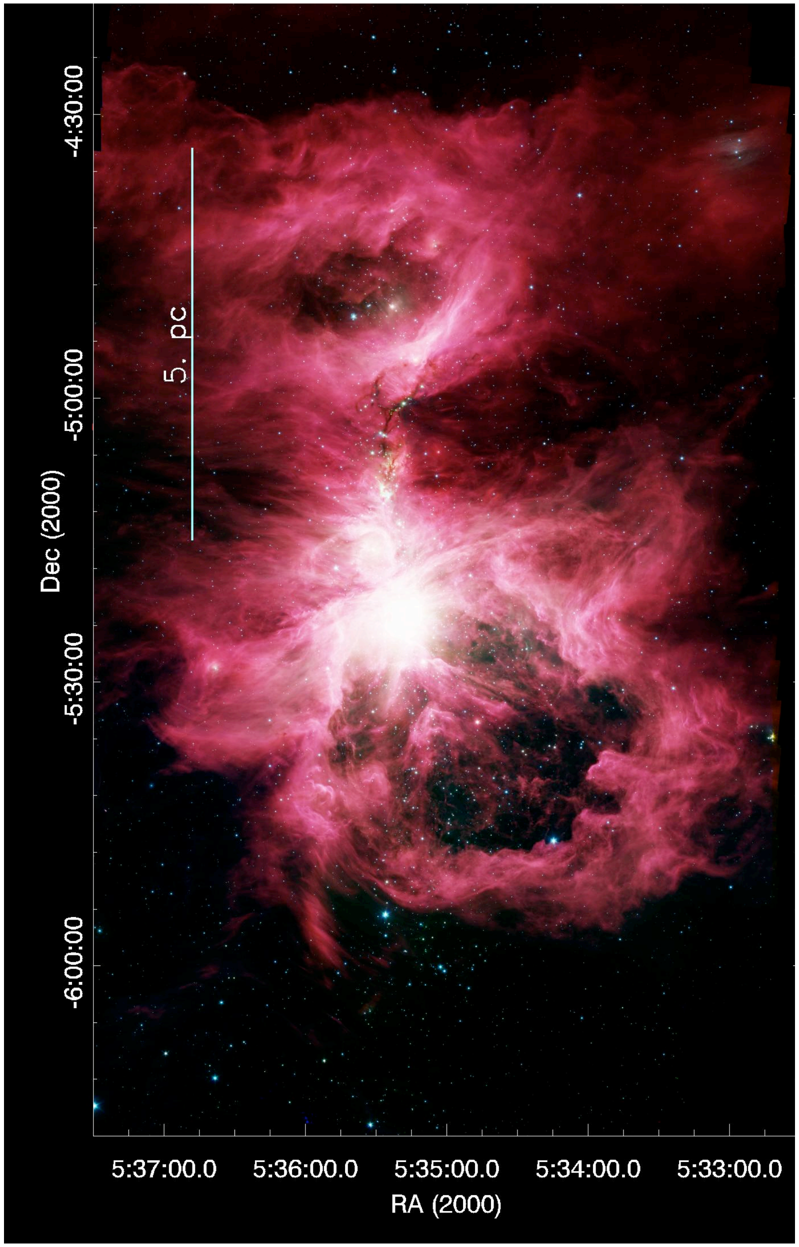
⇒ $(GM/r)^{1/2} \sim 5$ km/s; roughly what is observed (see next slide

- sound speed @ 10 K ~ 0.2 km/s ⇒ no thermal support
- "turbulent" motions ?

Velocity fields not particularly random



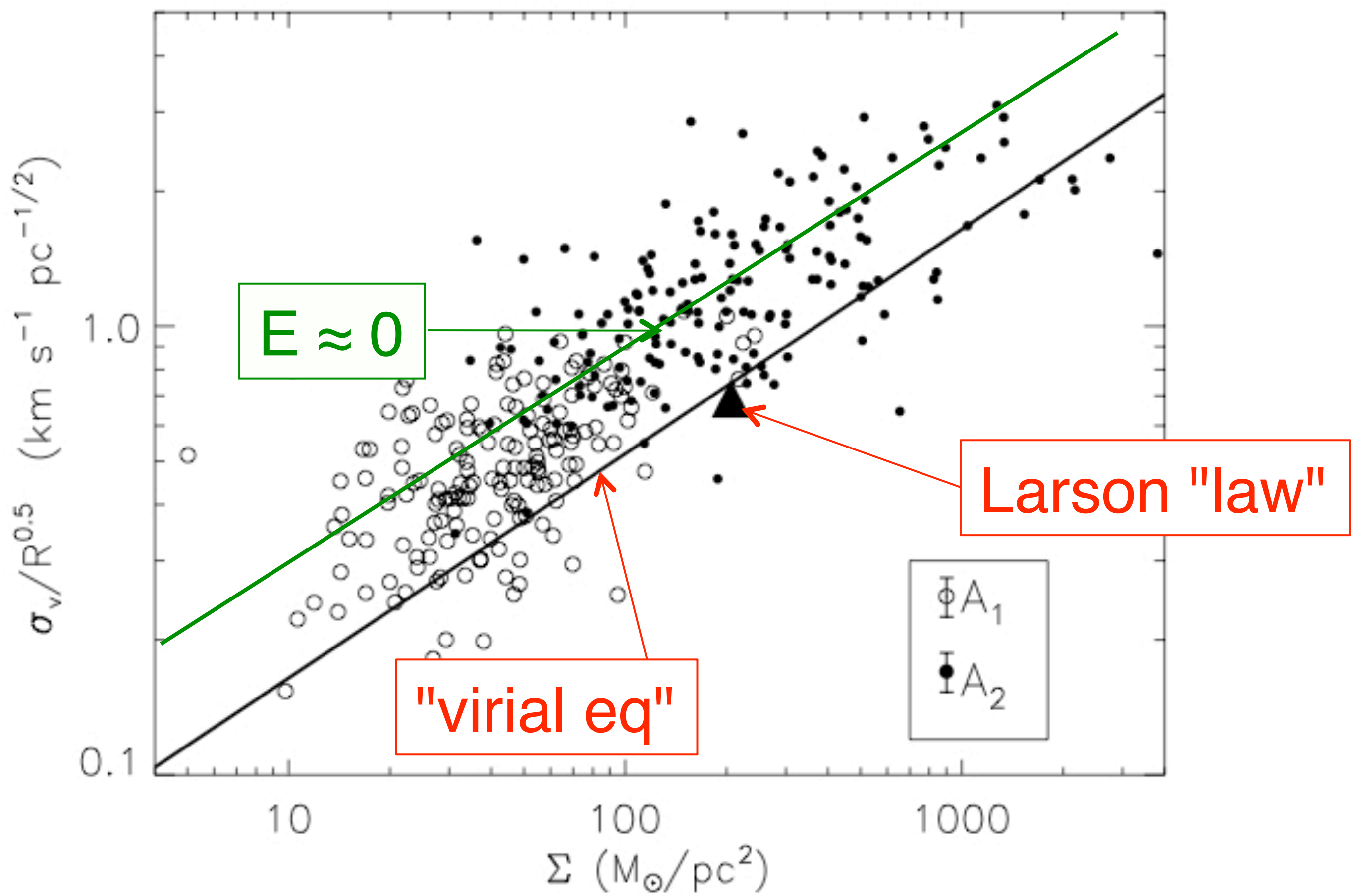
Orion A ^{13}CO; Bally et al. 1987



(N.B.; considering dense gas not being blown out by stellar winds etc.)

Molecular clouds

"virial" cloud masses; $\sigma_v \approx (G\,\Sigma\,R)^{1/2} \approx (GM/R)^{1/2}$



Heyer et al. 2009

Cloud masses are often estimated from a "virial equilibrium" assumption;

$$G\,M\,/R \sim v^2$$

Where does the "turbulence" come from to balance gravity?

If however motions are DRIVEN by gravity:

$$(2)\ G\,M\,/R \sim v^2$$

there is no mystery.

Difficult to avoid non-linear spatial acceleration by gravity in a complex geometry (neglect thermal etc. pressure)

Note: there IS non-gravitational turbulence… but supersonic velocities much easier to undertand as G

Free-fall timescale and homologous vs. non-linear collapse

Consider a sphere of radius R with uniform density ρ. At large scales, where gas pressure is not important, all radii fall in at the same time. One can see this from dimensional analysis:

$$v^2 \sim GM/R \sim G\rho R^2 ; \qquad \textit{thus} \qquad t_{ff} \sim R/v \sim (G\rho)^{1/2} . \qquad (1)$$

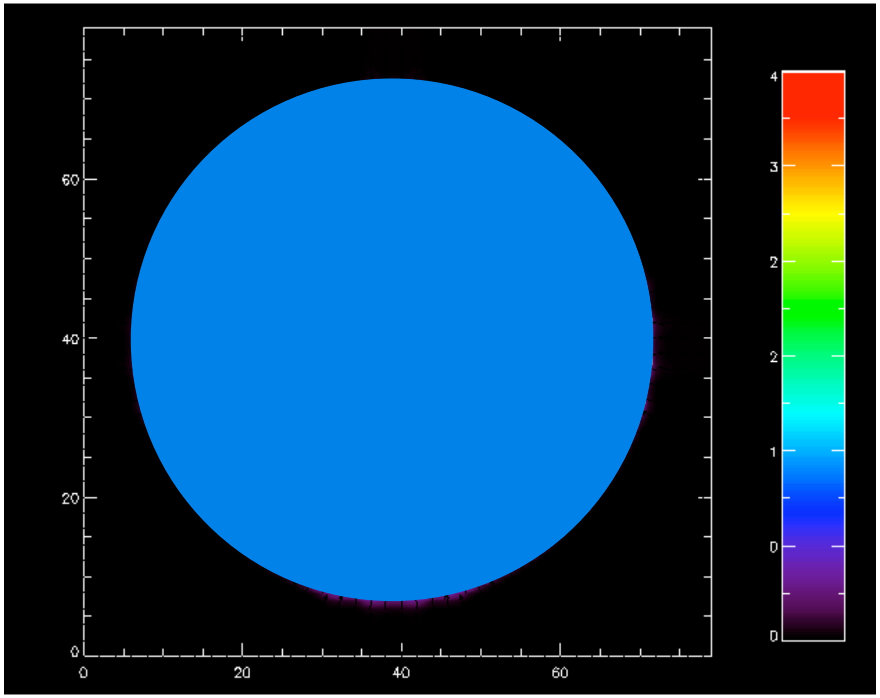
Molecular clouds cannot be supported by thermal pressure; (neglecting magnetic fields) global collapse is likely to occur.

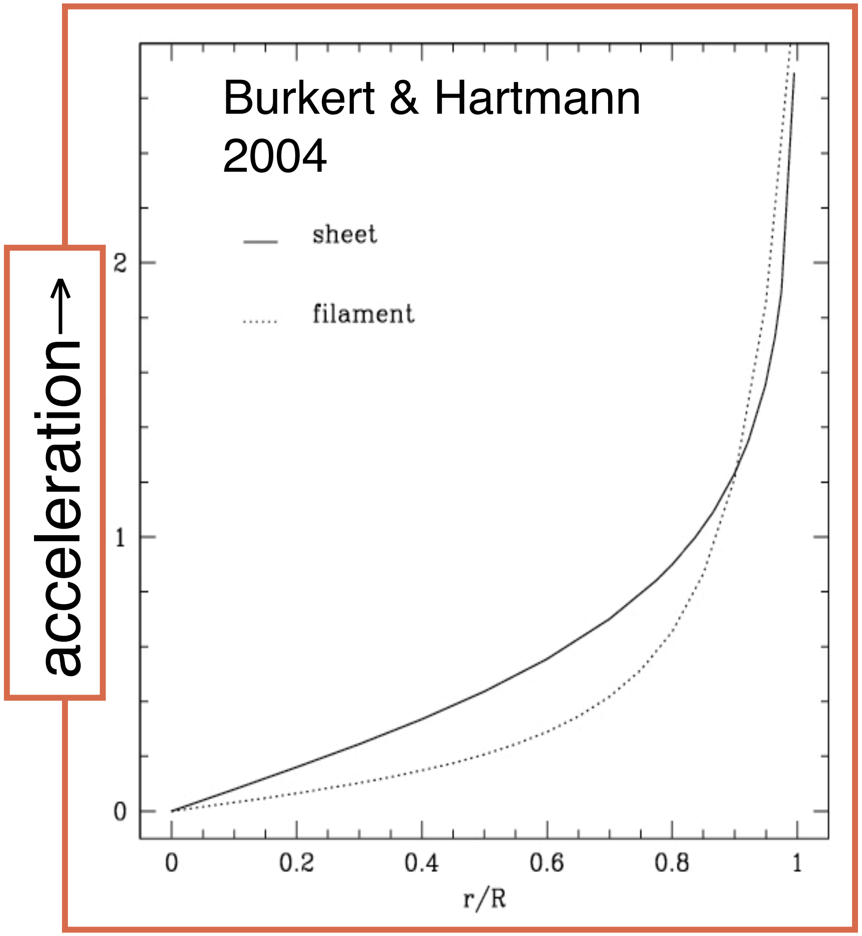
However, the global nature of gravity in a non-uniform, non-spherical volume results in non-linear accelerations as a function of position, so that it is essentially impossible to prevent collapse somewhere (unless gravity is too weak or expansion is too large).

Finite sheet evolution with gravity

uniform surface density Σ, isothermal, circular sheet:

⇒ pileup of material!

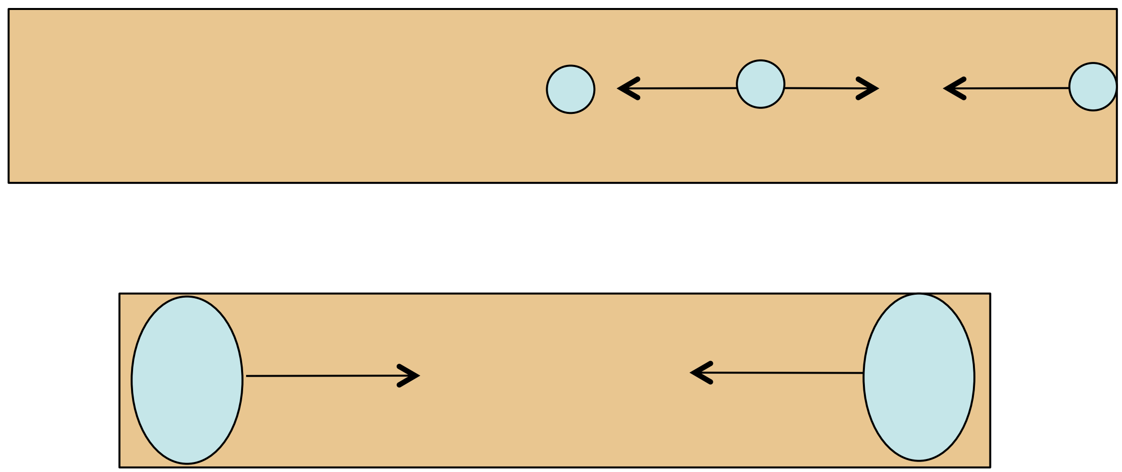




non-linear gravitational acceleration vs. position

⇒ make filaments

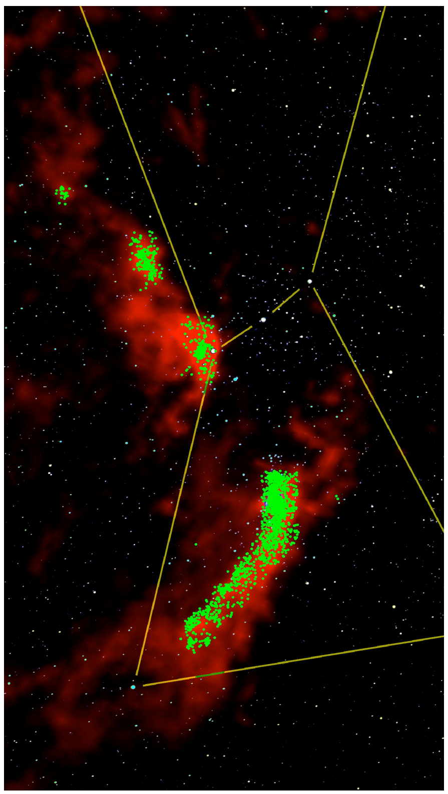
Non-linear acceleration vs. position in filaments (Bonnell et al 1992)- “edge effect” in ~ 1 dimension



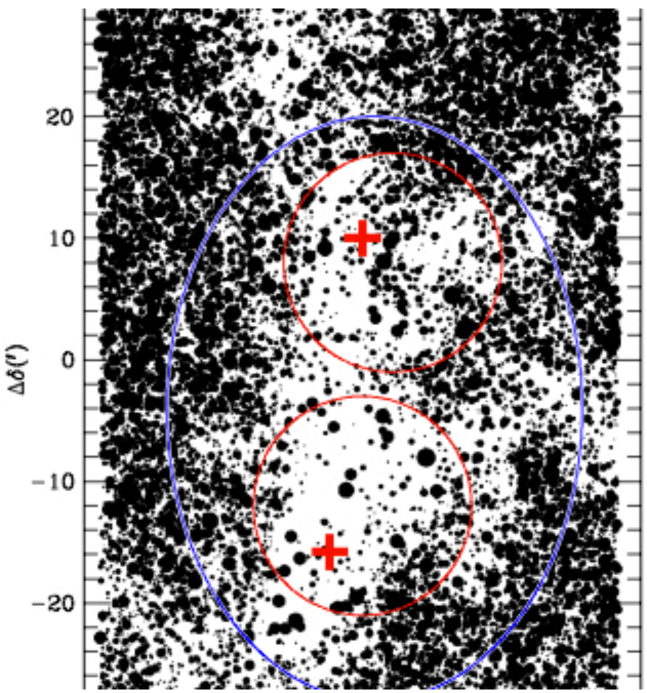
make concentrations (clusters) at ends

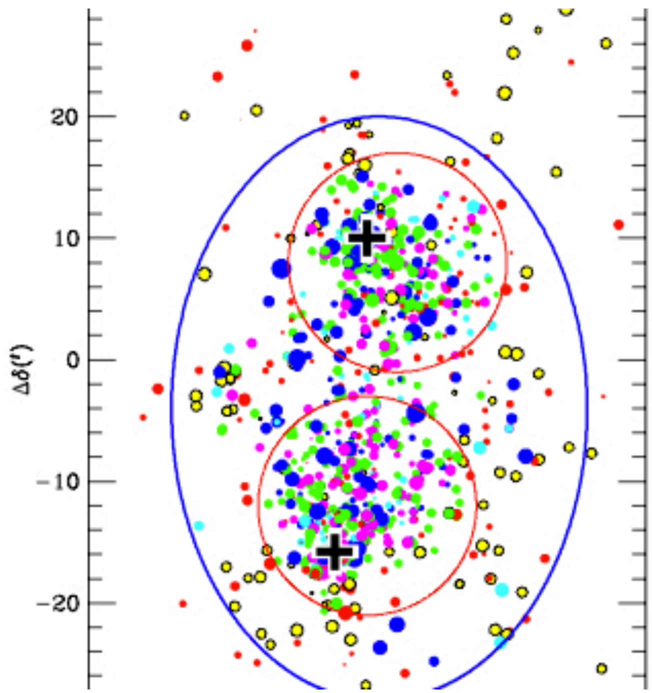
Other evidence for gravitational collapse? Making clusters near ends of ~ filaments

Orion A; Megeath

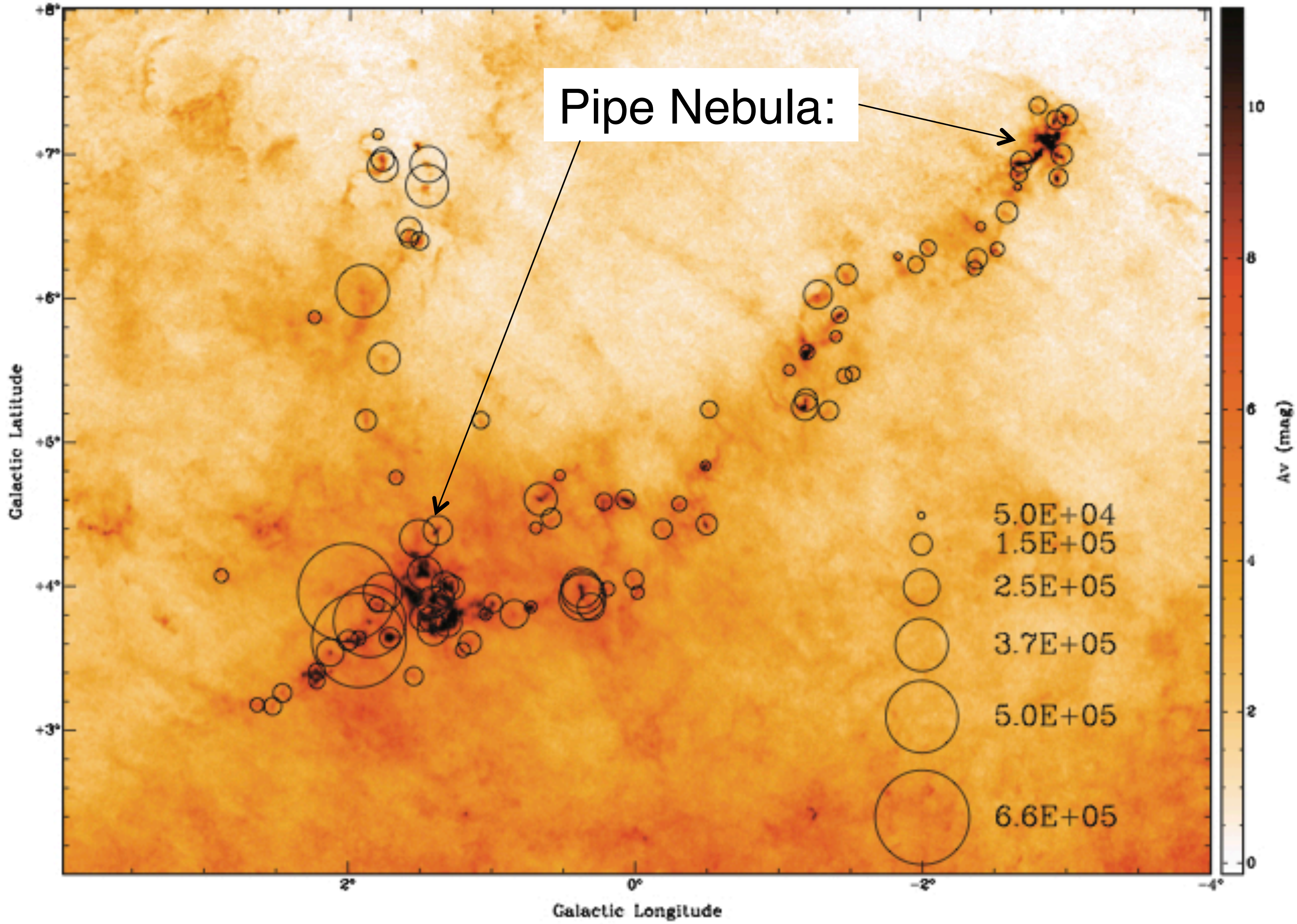


NGC 2264; Sung + 2008



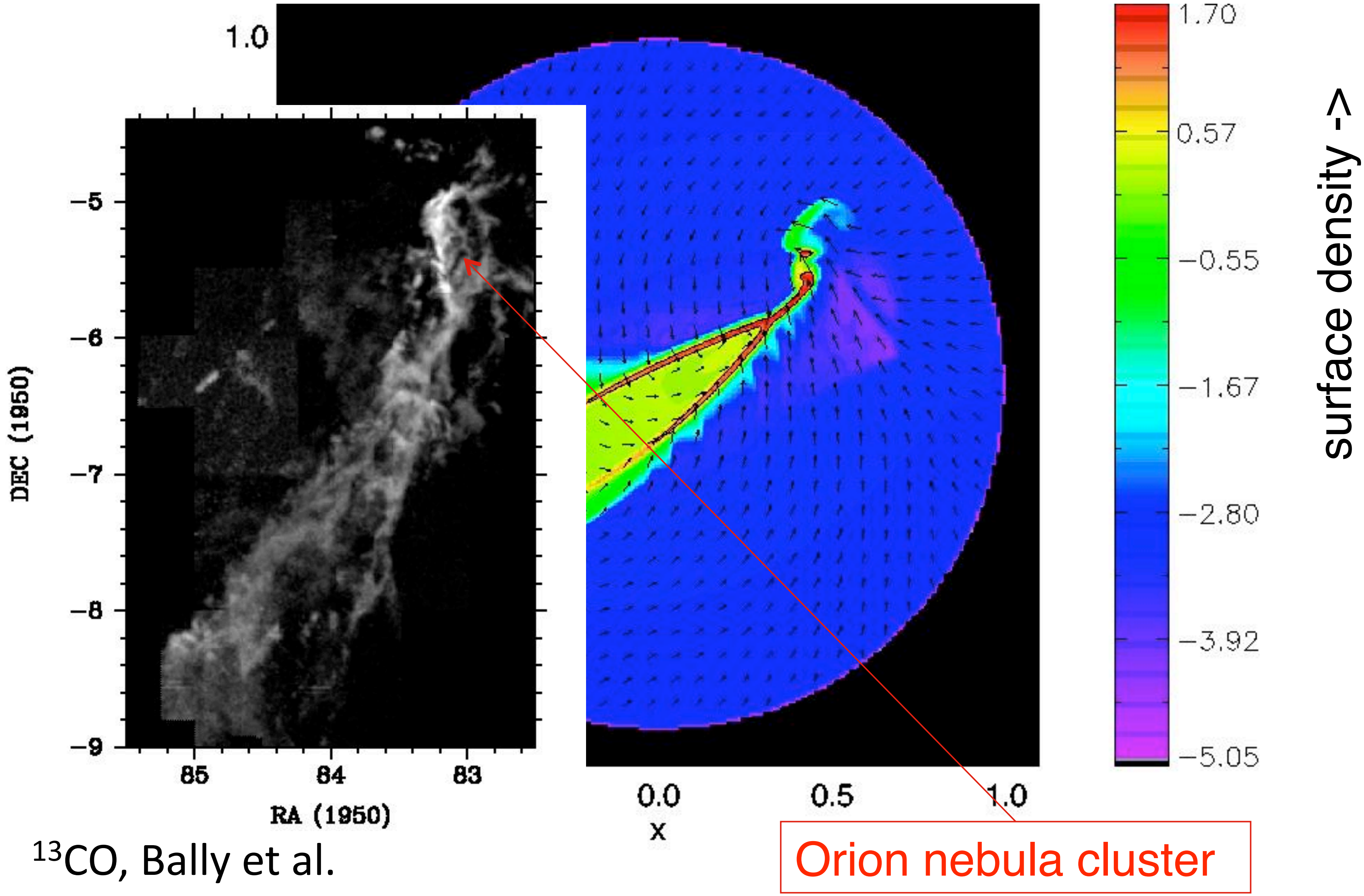


extinction



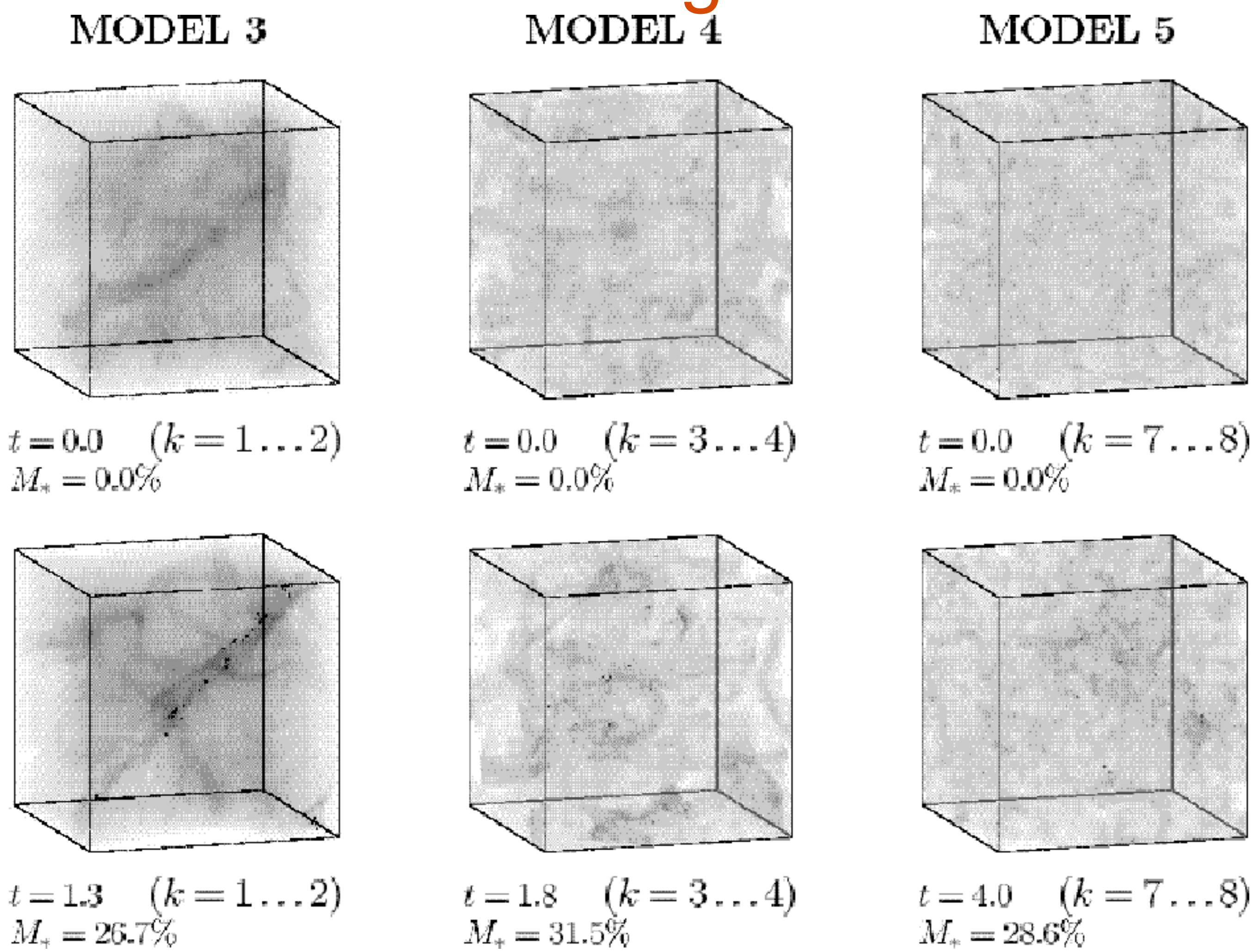
Lada et al. 2008

“Orion A” model; gravitational collapse of finite, massive, elliptical, rotating sheet (Hartmann & Burkert 2007)



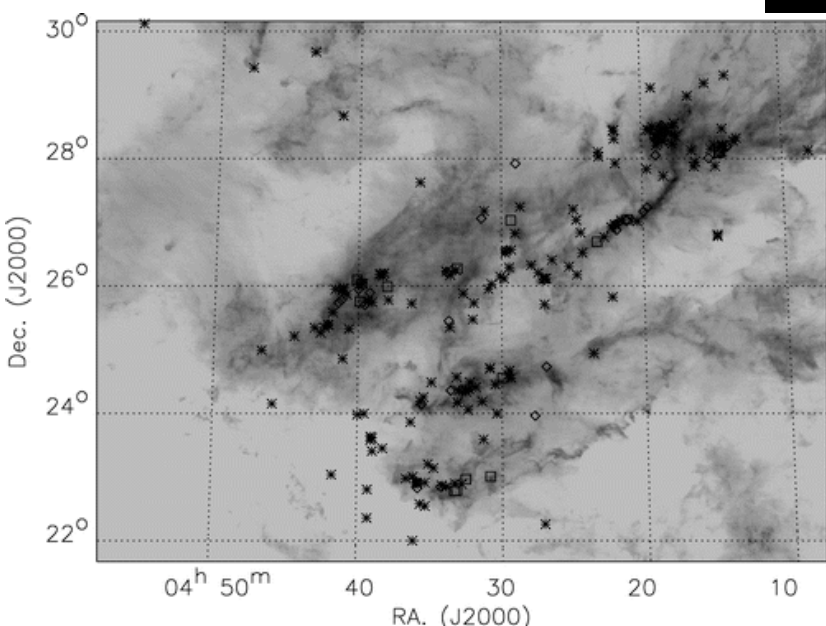
^{13}CO, Bally et al.

“turbulent driving” can also make filaments if large-scale

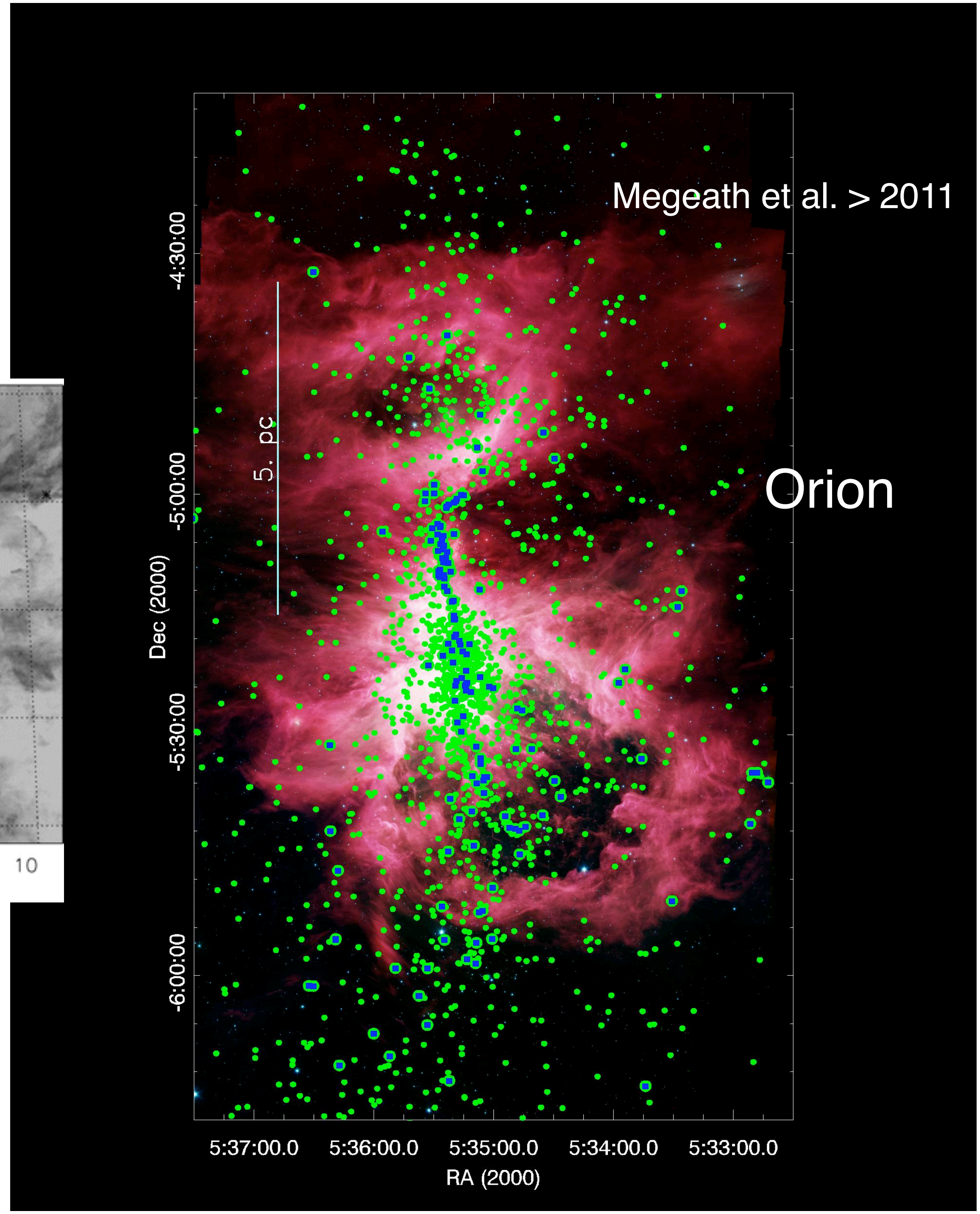


Klessen 2001, Klessen, Heitsch & Mac Low 2000; Krumholz, Matzner, Klein, McKee… low wavenumber driving ⇒ extended filaments;

gravitational collapse good for making filaments, where many stars form (see lecture 2)

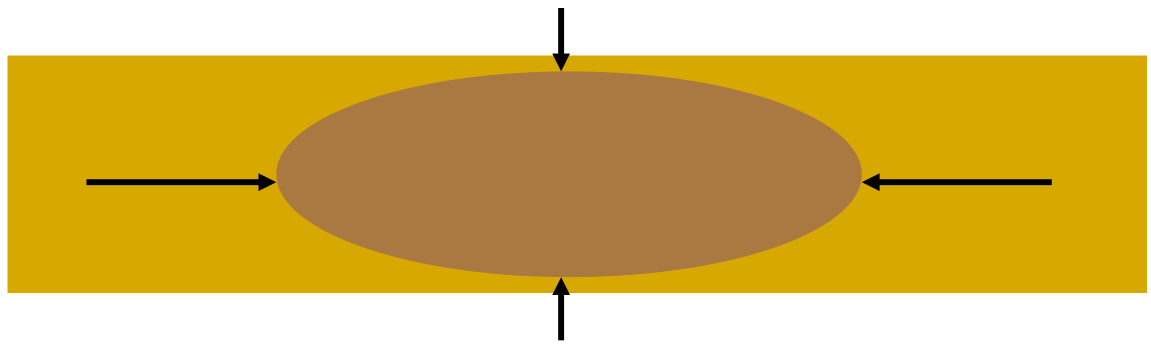


Taurus

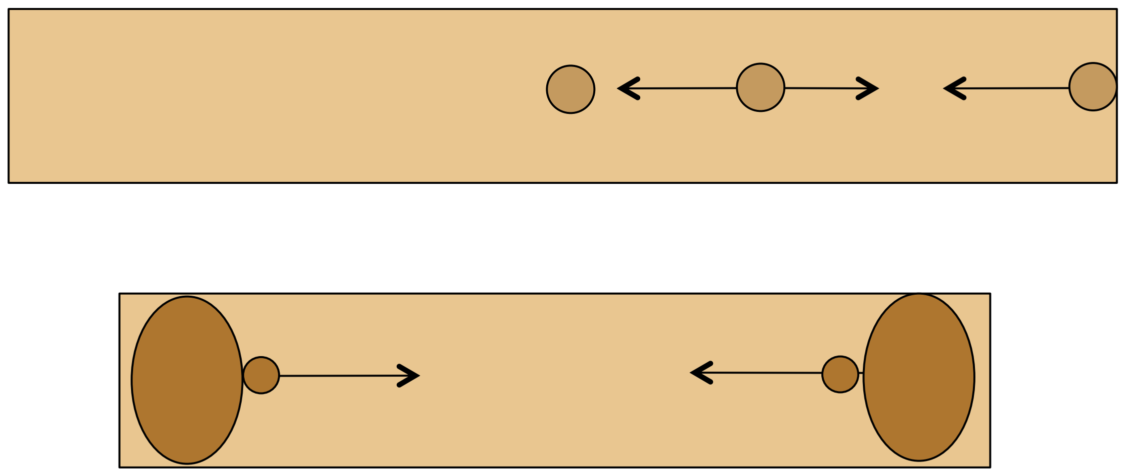


protostellar clouds ("cores") in filaments?

Fragment gravitationally in elongated structures along filaments (e.g., Larson 1985)?



$\lambda_J > H$ (because H is length scale pressure can support against gravity) simple dispersion relation doesn’t hold in finite filaments; on-linear acceleration vs. position (Bonnell et al 1992)



Problem: for finite filaments, competition from global vs. local collapse.. need DENSE “seeds” which cool, become denser, and collapse faster than the global infall (could use some numerical investigations…)

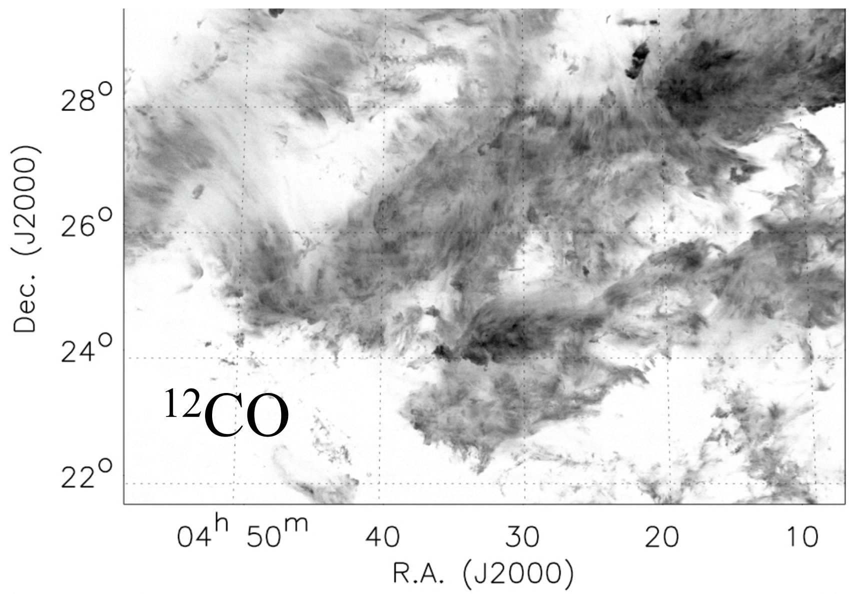
we want to make stars, masses << cloud mass.

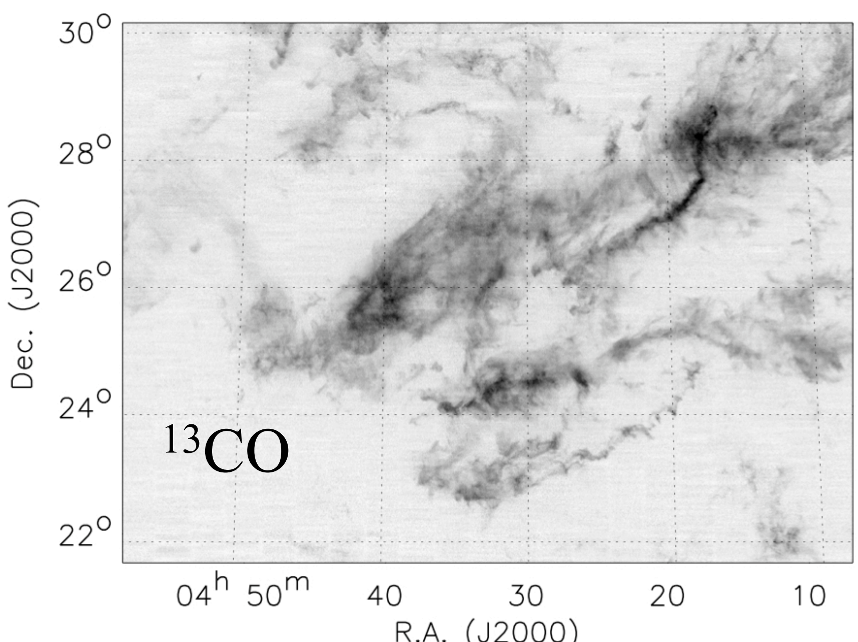
how? Must make dense “seeds” which can collapse faster than possible global collapse, even in filaments.

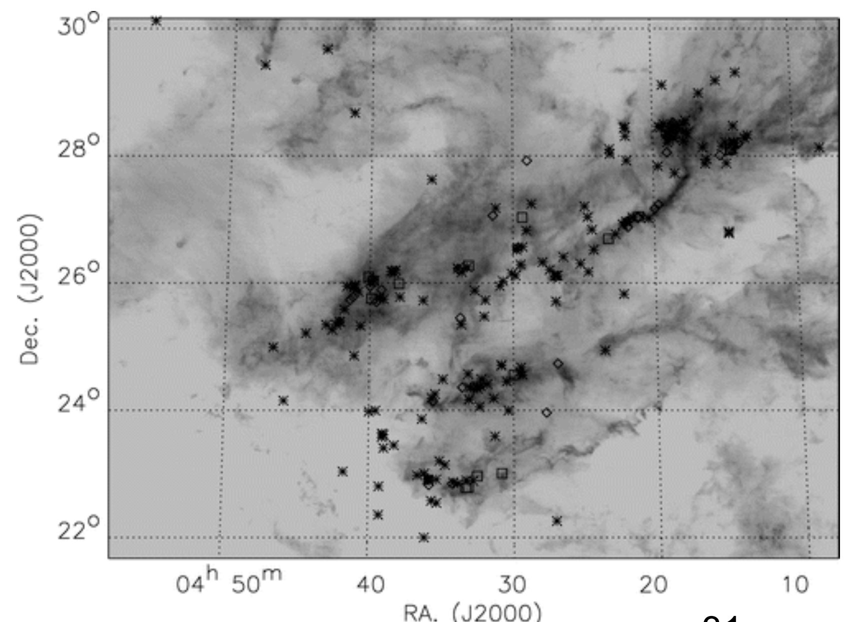
Turbulence, amplified by gravity.

star-forming molecular clouds DO have "turbulent" substructure – this is needed to break clouds into much smaller pieces (stars)…

Expect turbulence because "Reynolds number" is large;
Re ~ u d / ν >> 1



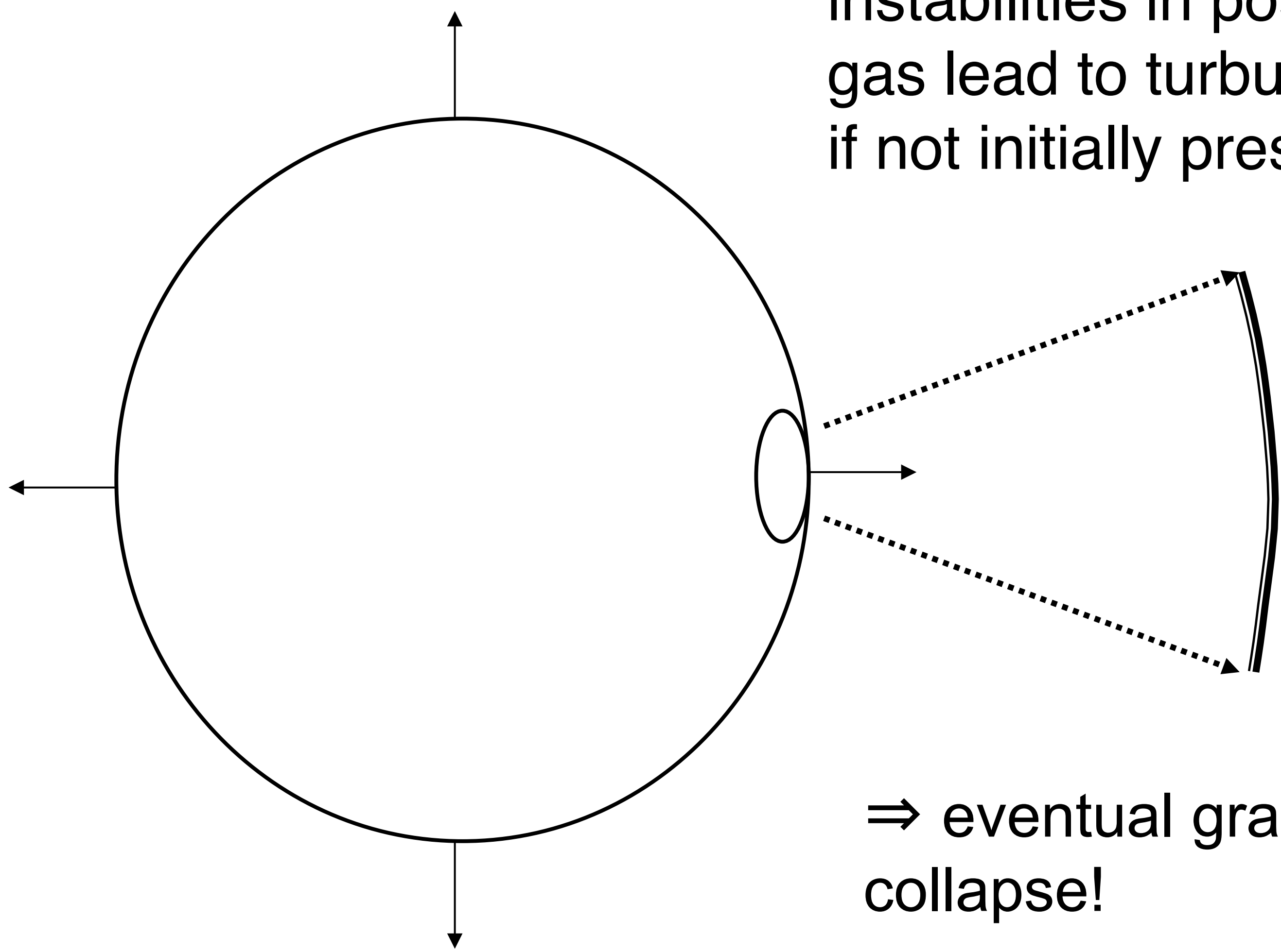




Taurus: Goldsmith et al. 2008

Finite sheet evolution with gravity

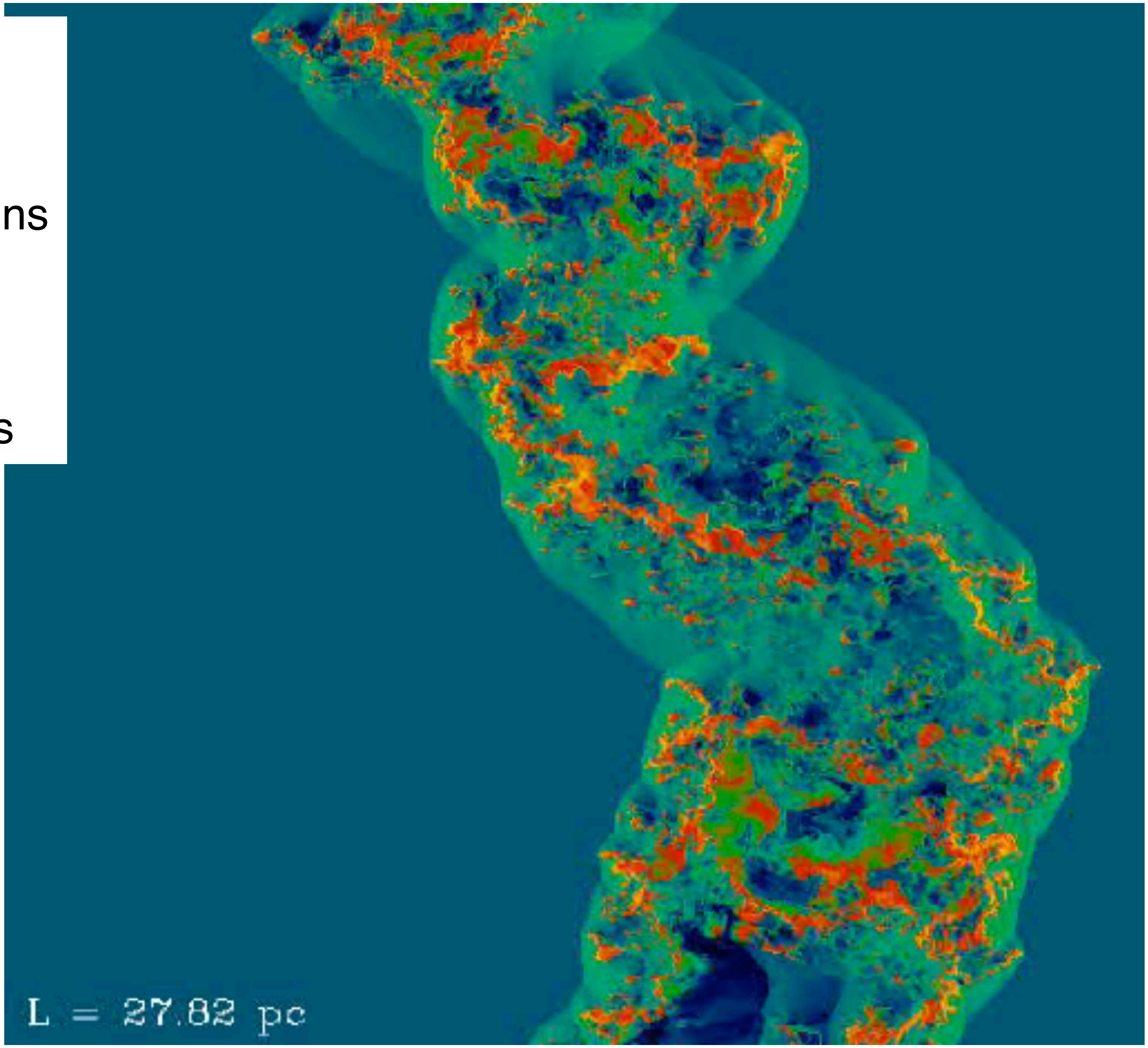
Let's simplify: piece of bubble wall ≈ sheet



instabilities in post-shock gas lead to turbulence even if not initially present

⇒ eventual gravitational collapse!

inflow
without
gravity;
perturbations
make
turbulent
density
fluctuations

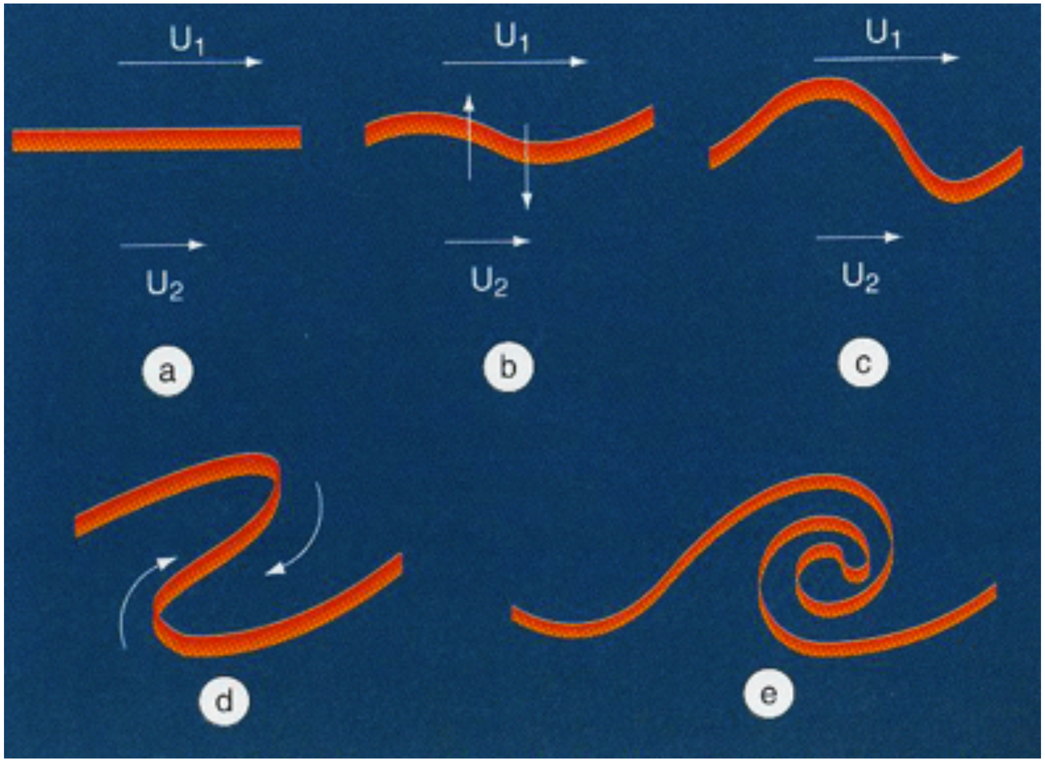


uniform inflow, perturbed interface

Heitsch, 2007 (2D Kelvin-Helmholtz)

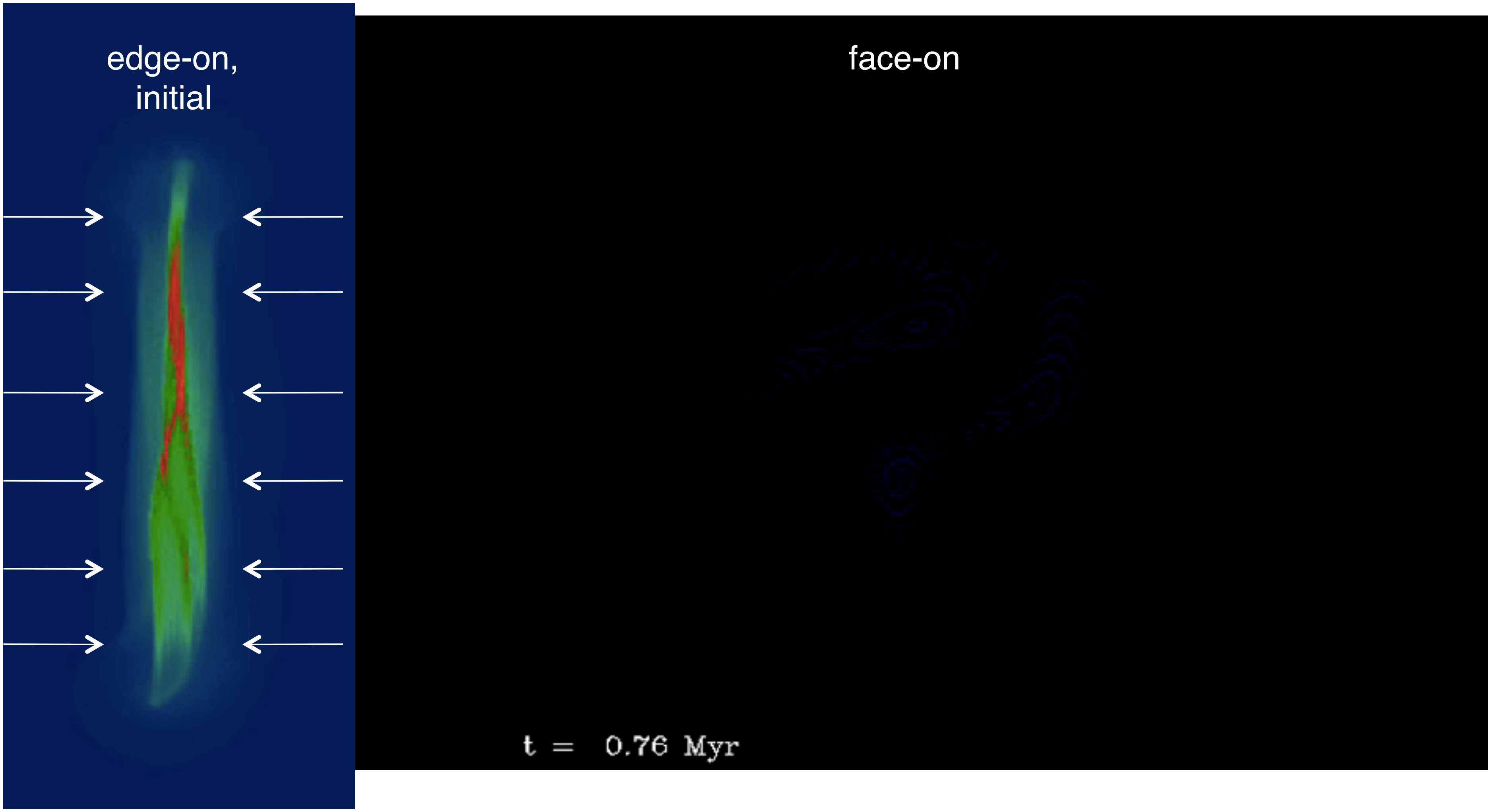
Kelvin-Helmholtz instability

(start with Bernoulli effect on wave interface)

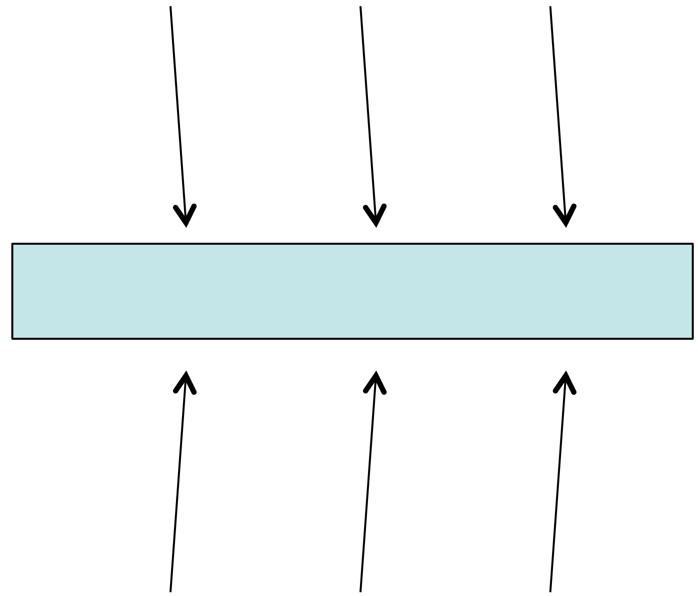


http://hmf.enseeiht.fr/travaux/

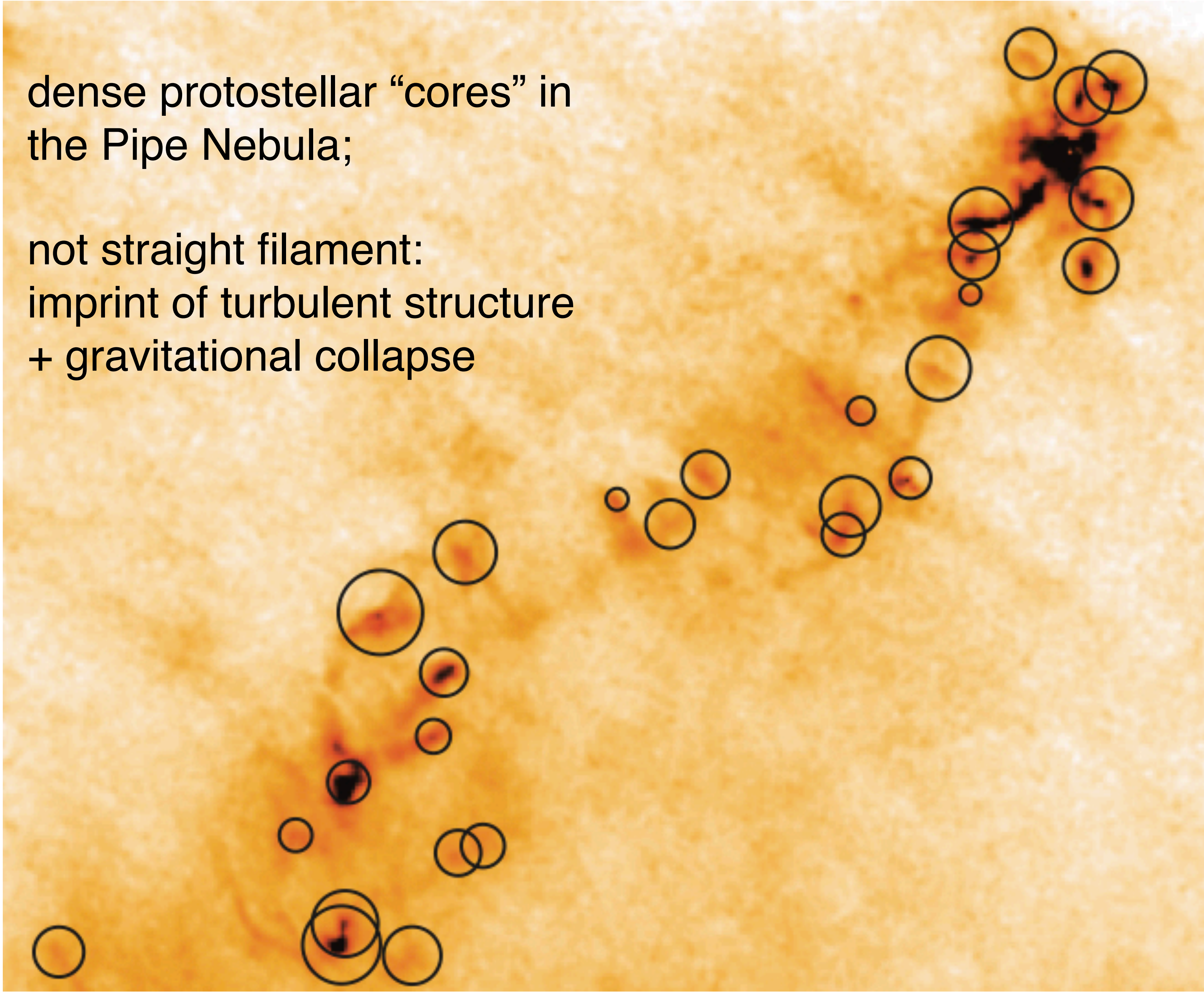
Now, let's add gravity to the turbulence initially generated at the shock interface… fragments in filaments!



Heitsch+ 2007, 2008; Hennebelle, many; Vazquez-Semadeni+ 2007, 2010



gravitational collapse to a filament, subsonic region can fragment (Gong & Ostriker, Ballesteros-Paredes et al.)


dense protostellar “cores” in
the Pipe Nebula;
not straight filament:
imprint of turbulent structure
+ gravitational collapse

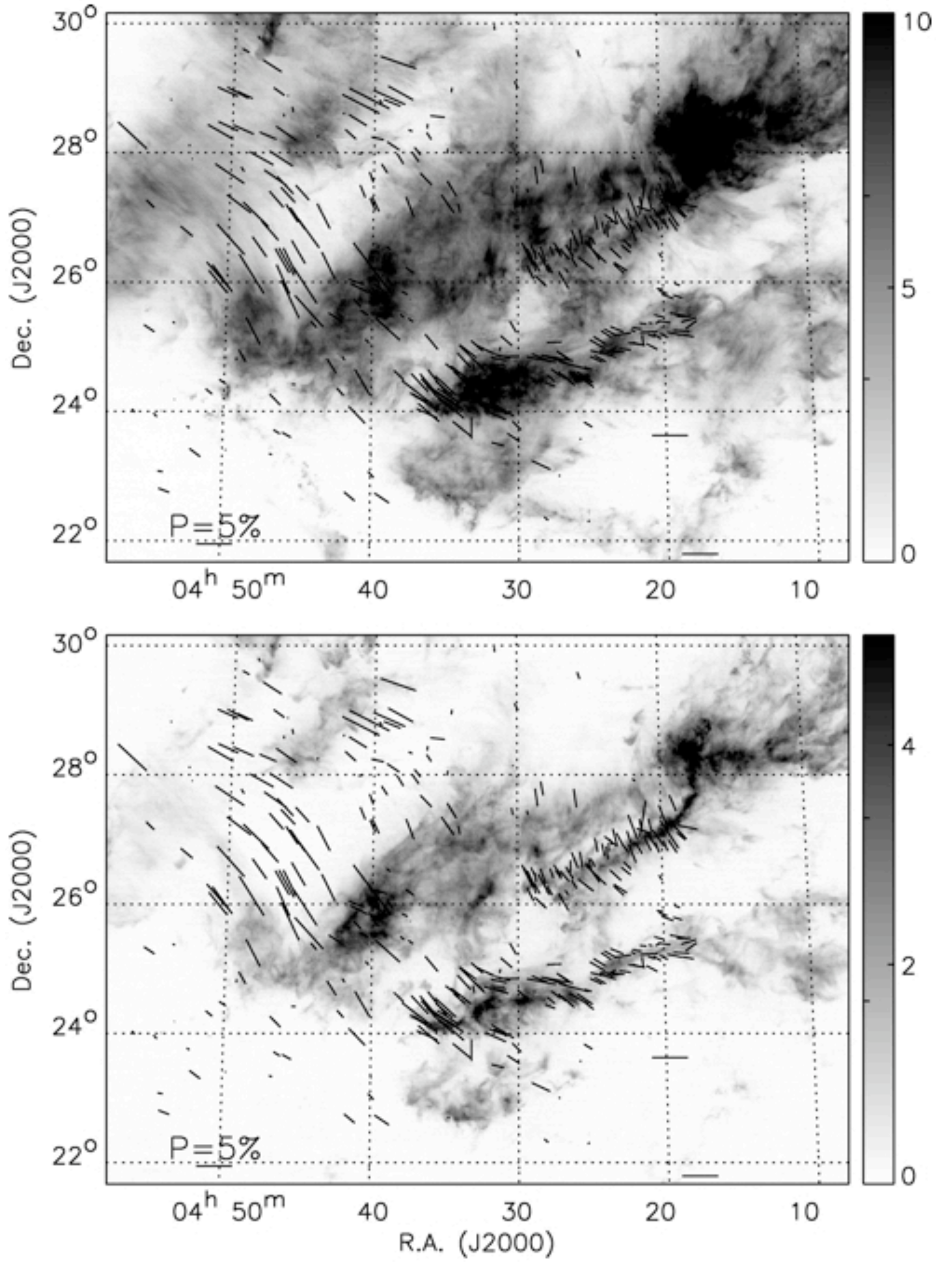
Can we avoid global gravitational collapse with magnetic fields – turbulent or otherwise?

(local collapse?)

Magnetic fields are present; can be dynamically significant

Does magnetic field control flow or does the gas carry the field along with it?

High and low density regions; B strong/weak?



Optical polarization of background stars (lines are ll to observed planes of polarization)

Energy balance principle, again

$(B^2/8\pi)$ x volume $\approx GM^2/R$

$\Rightarrow B^2 R^4 \approx GM^2 \Rightarrow B R^2 \approx G^{1/2} M$

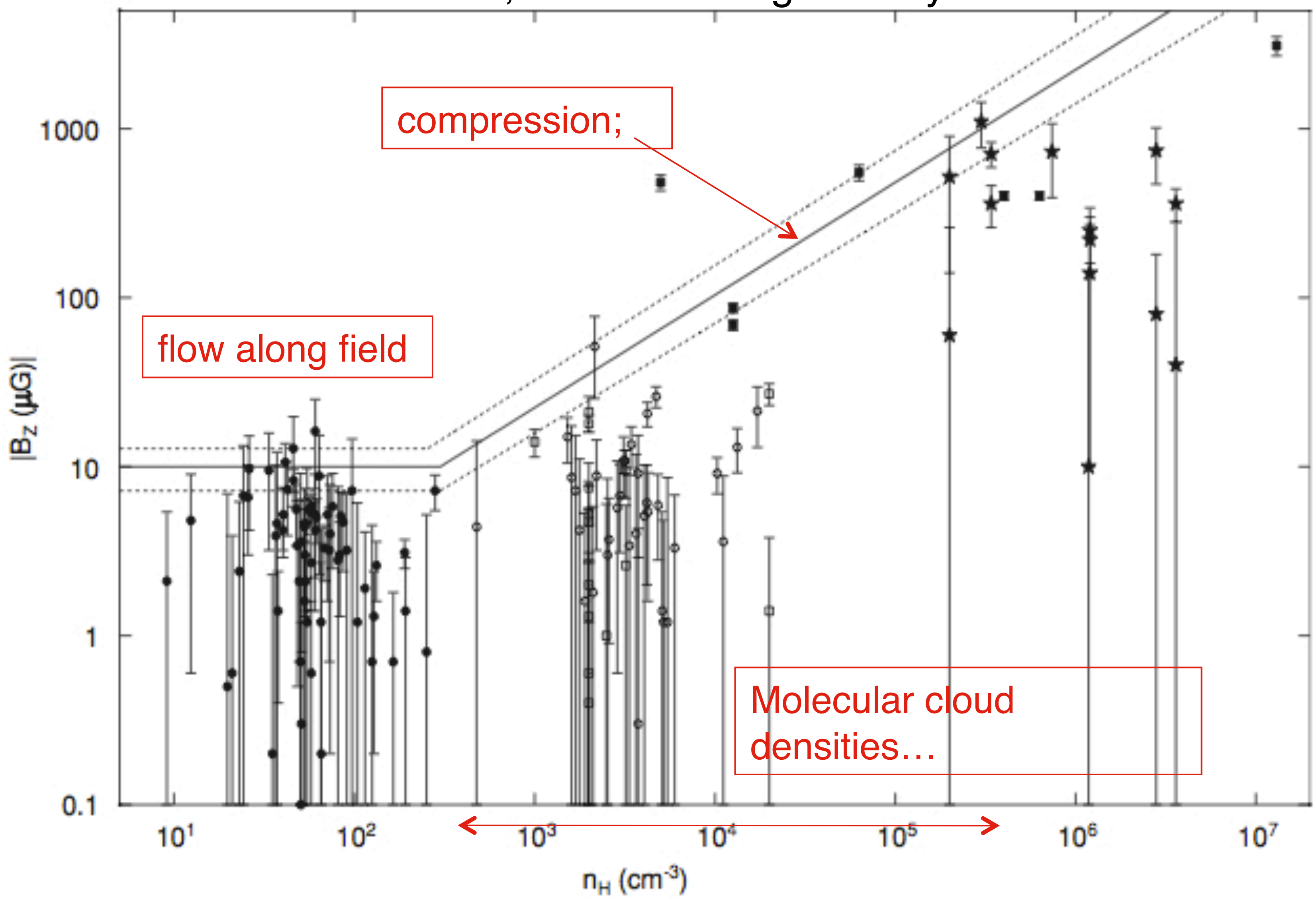
(magnetic flux through cloud vs. M);

if M > critical flux, can collapse;
if M < critical, cannot collapse unless
gas can slip through field lines

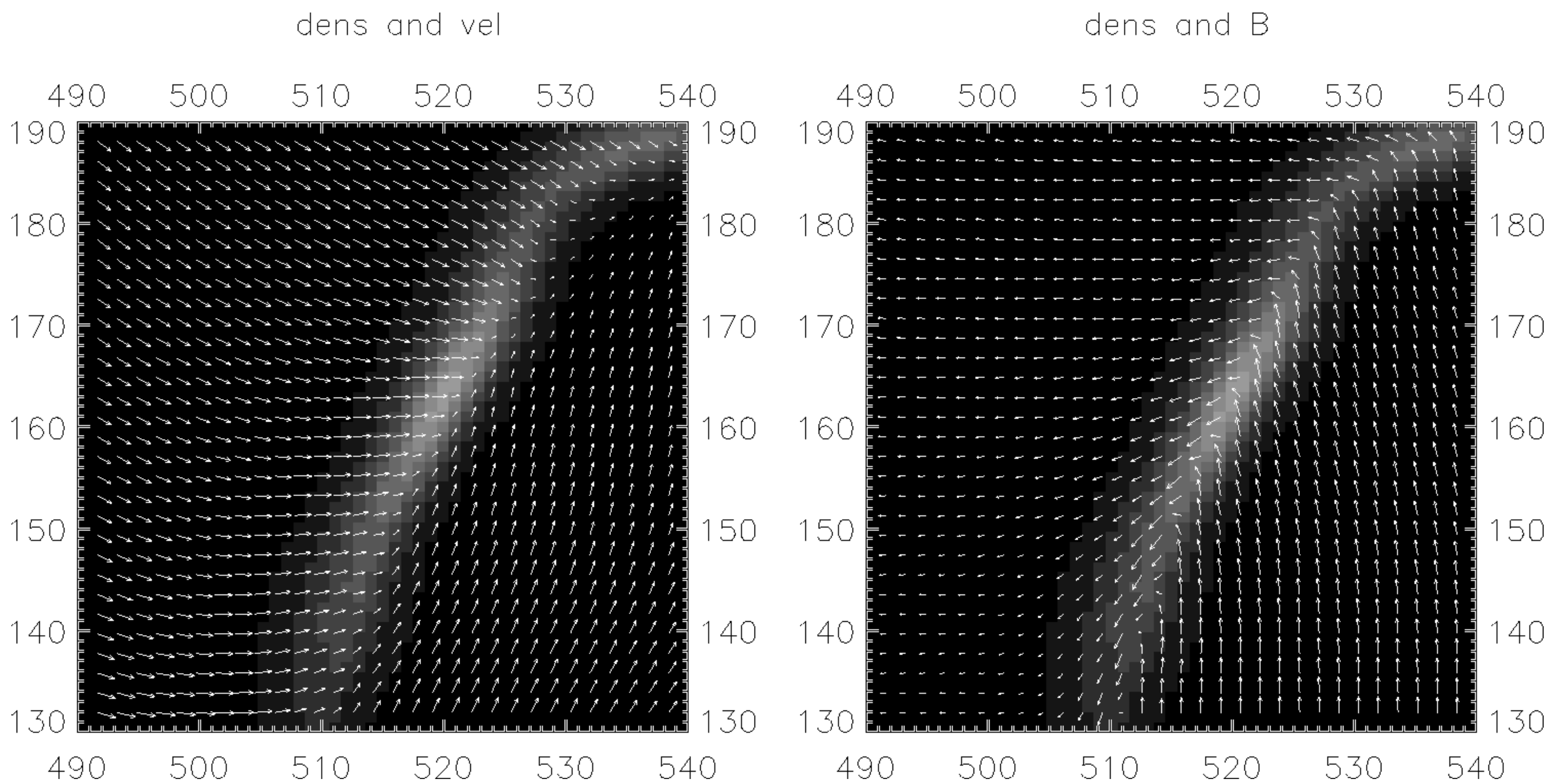
other form: if CLOSED mass;

$B R^2 \propto \rho R^3 \Rightarrow B \propto \rho^{2/3}$

Crutcher et al. 2010, ApJ 725, 466: infer wide variation in mass-to-flux ratios in dense clouds; some are magnetically-weak



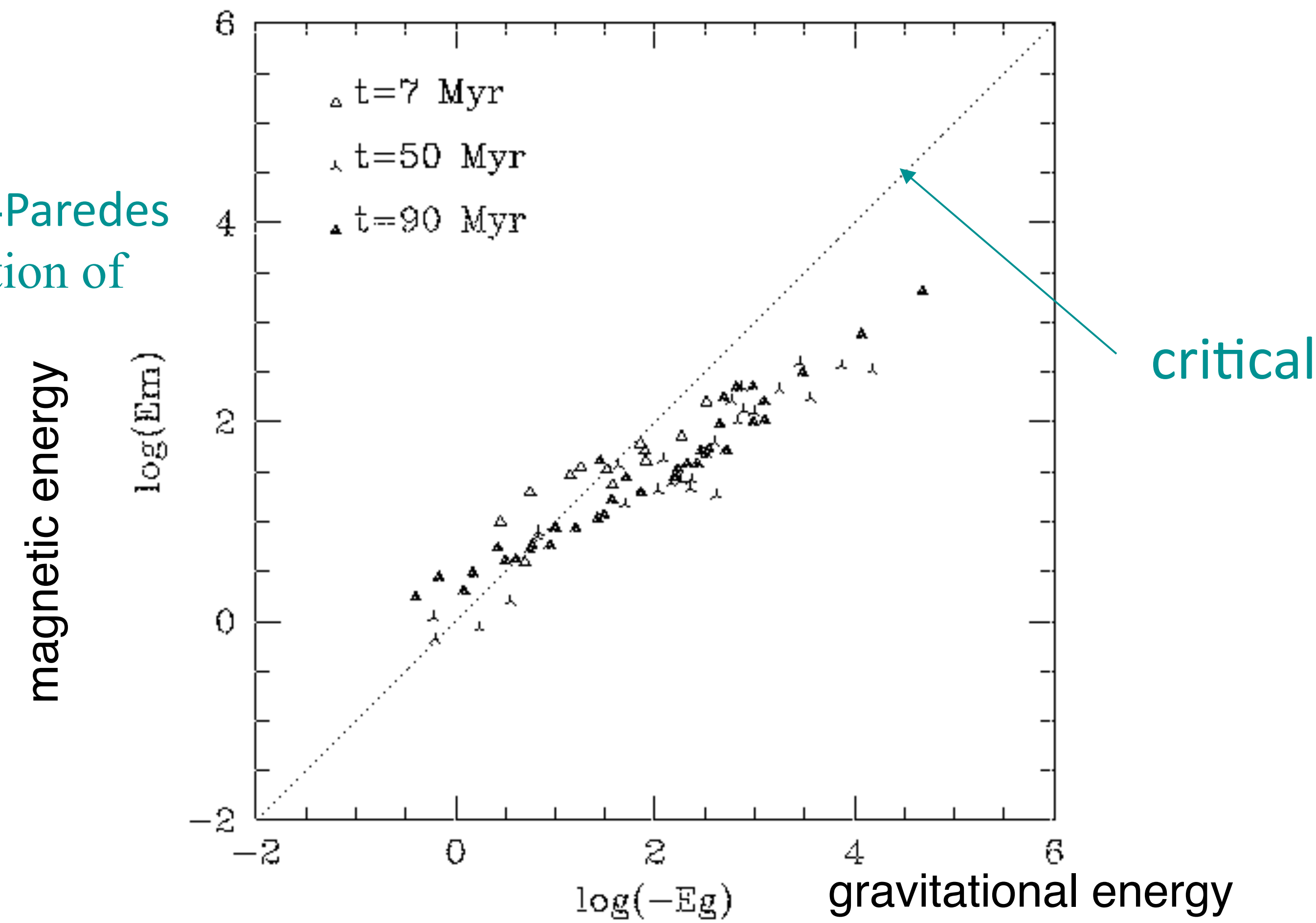
Cloud formation starting with large scales; flow ALONG B; low-density regions are magnetically strong, mass piles up until magnetically weak. because otherwise won't pile up



Ballesteros-Paredes, Hartmann, & Vazquez-Semadeni 1999

Cloud formation; supercritical at high density and vice versa

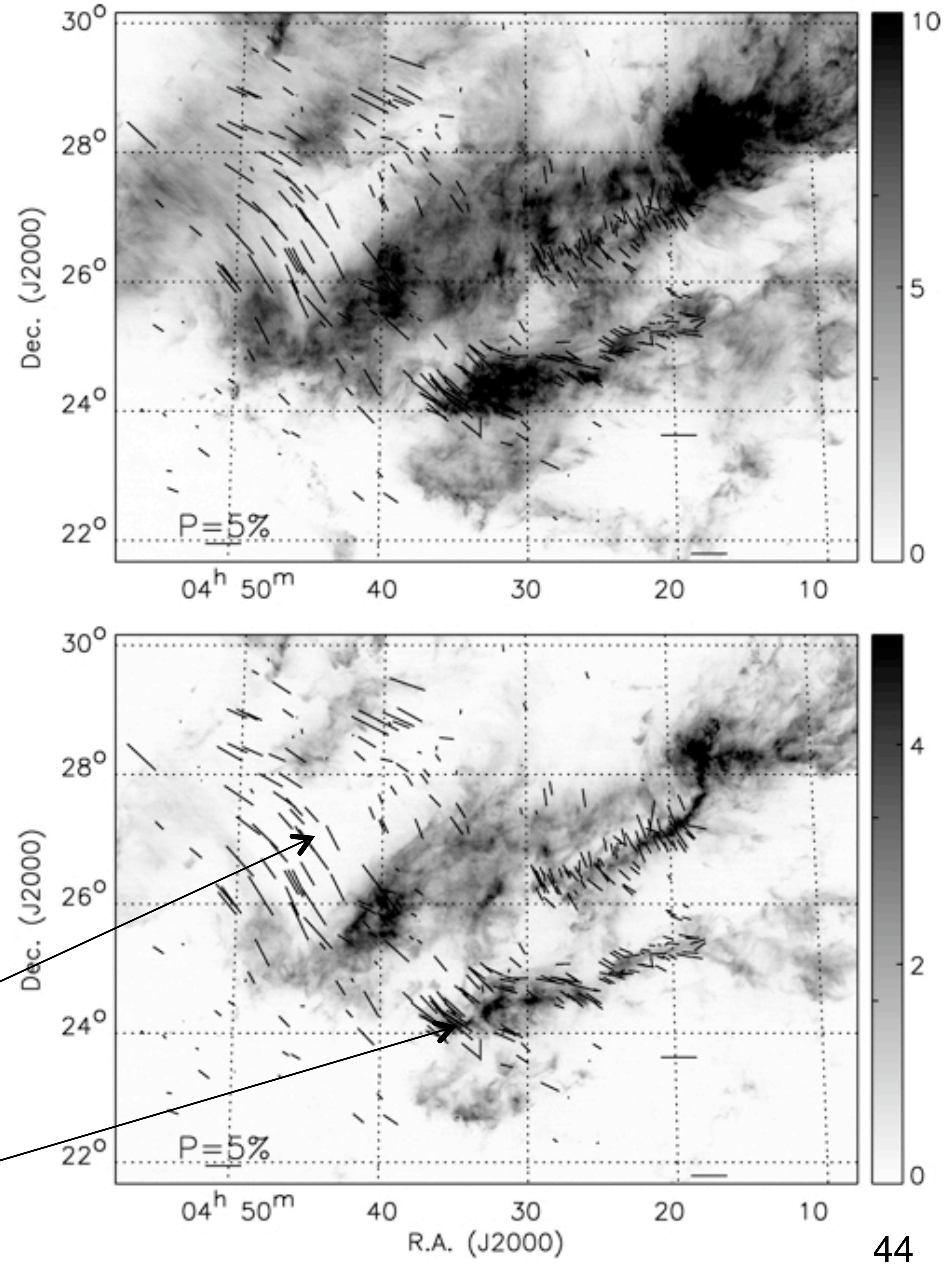
Ballesteros-Paredes 2-D simulation of clouds



Magnetic field controlled by pressure, eventually by gravity; Bertoldi & McKee; GMCs are globally supercritical

Magnetic support may be critical in reducing star-forming efficiency; i.e., much of the cloud could be magnetically-strong, while only densest parts B-weak (see Price & Bate 2009). Crutcher et al infer wide variation of B/ρ in dense clouds…

suggest collapse of individual cores NOT slowed by B



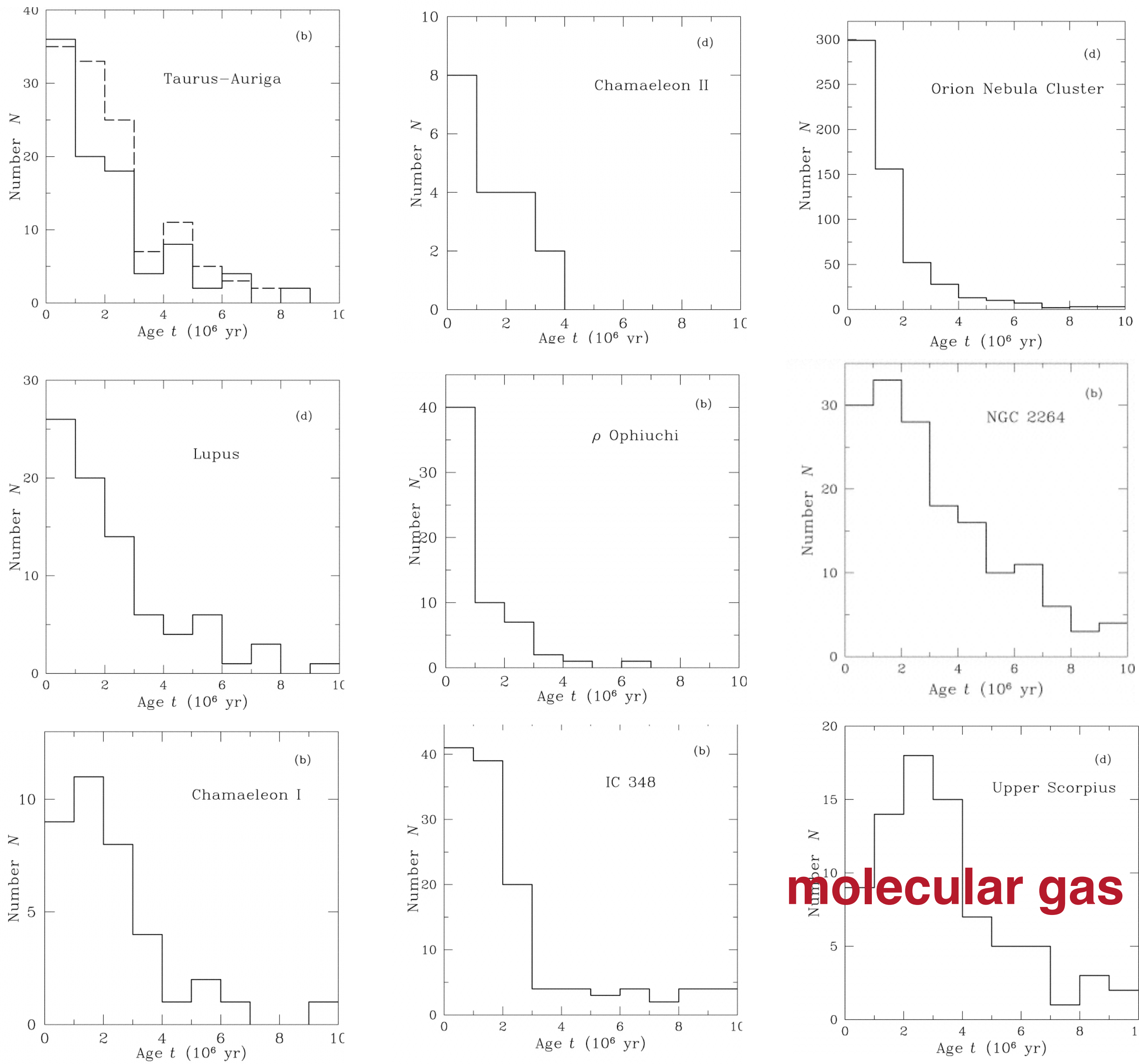
B strong?

B weak?

Magnetic turbulent support?

It was hoped that long-lived magnetic (Alfven) waves (non-compressive in first order) could support molecular clouds against their self-gravity

this appears not to work because the turbulence in the clouds is SUPERSONIC. In principle, A-waves can be supersonic while linear; however, given the tangled geometries of field lines and the non-perfect alignment of turbulence-generating energy input (i.e., stellar/disk winds, ionization fronts, SN,...) the waves will shock and dissipate rapidly, roughly on the order of a crossing time (Heitsch et al. 1999; Stone et al. 1998)



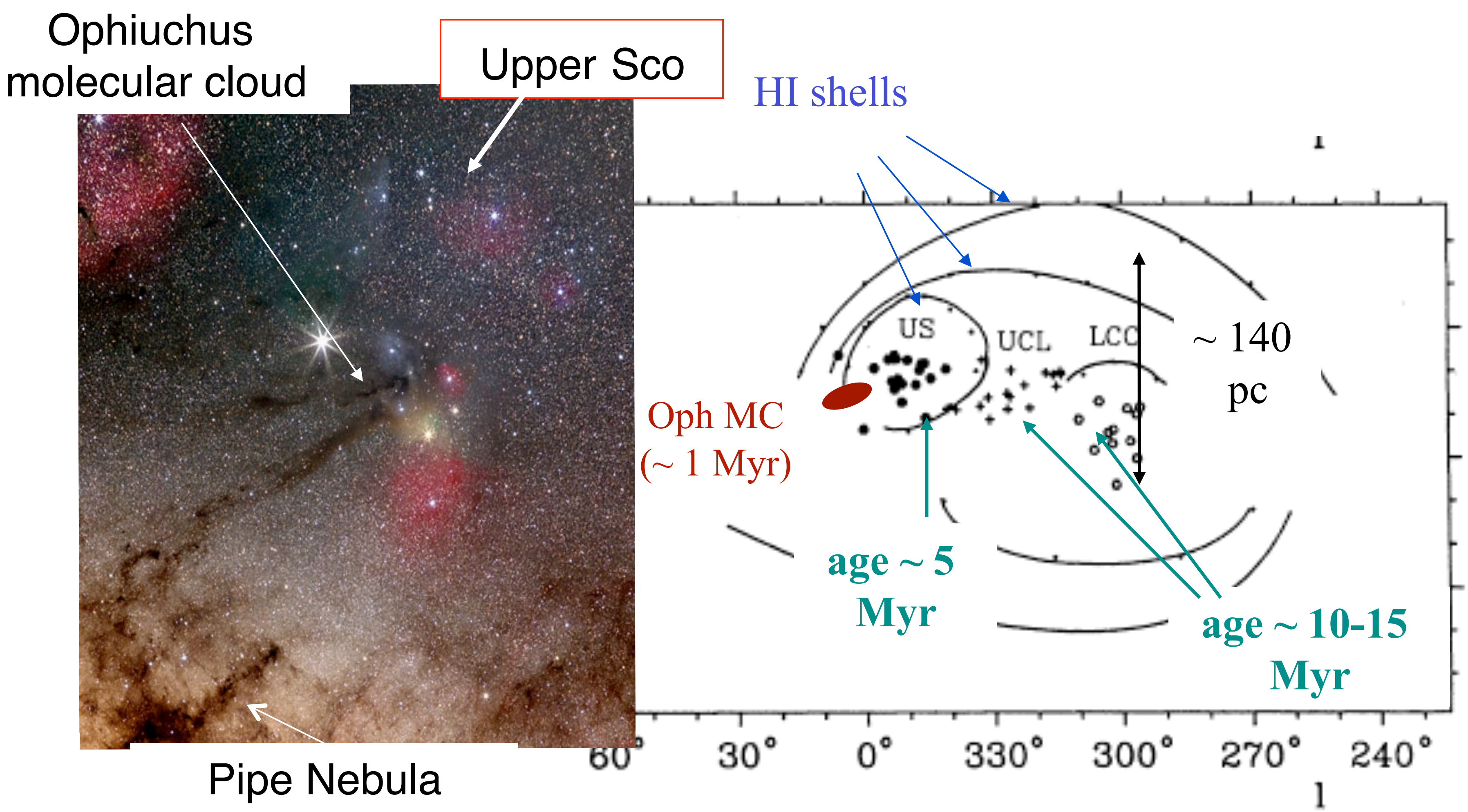
Palla & Stahler

star-forming regions; age peaks ~ 1-2 Myr ago; no slow formation; against magnetic slowing of collapse by ambipolar diffusion

molecular gas removed!

Sco OB2: quickly emptied by winds, SN

de Geus 1992; Preibisch & Zinnecker 99, 02



Supplemental material

Formation of H_2 on dust grains

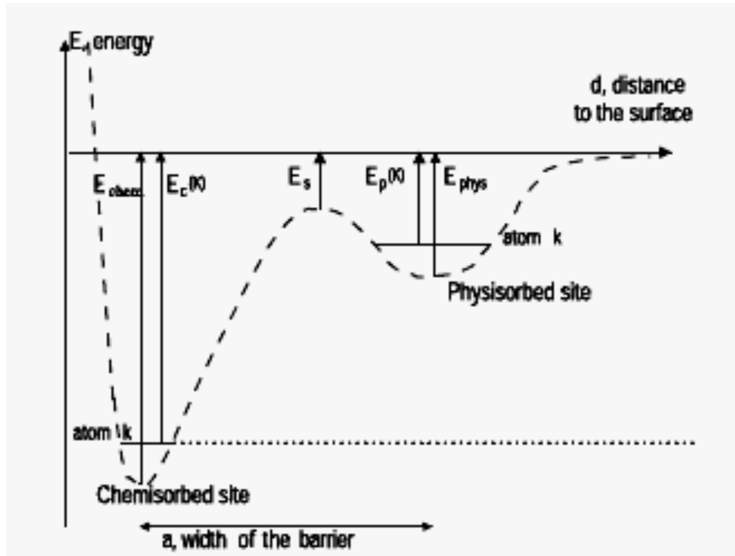


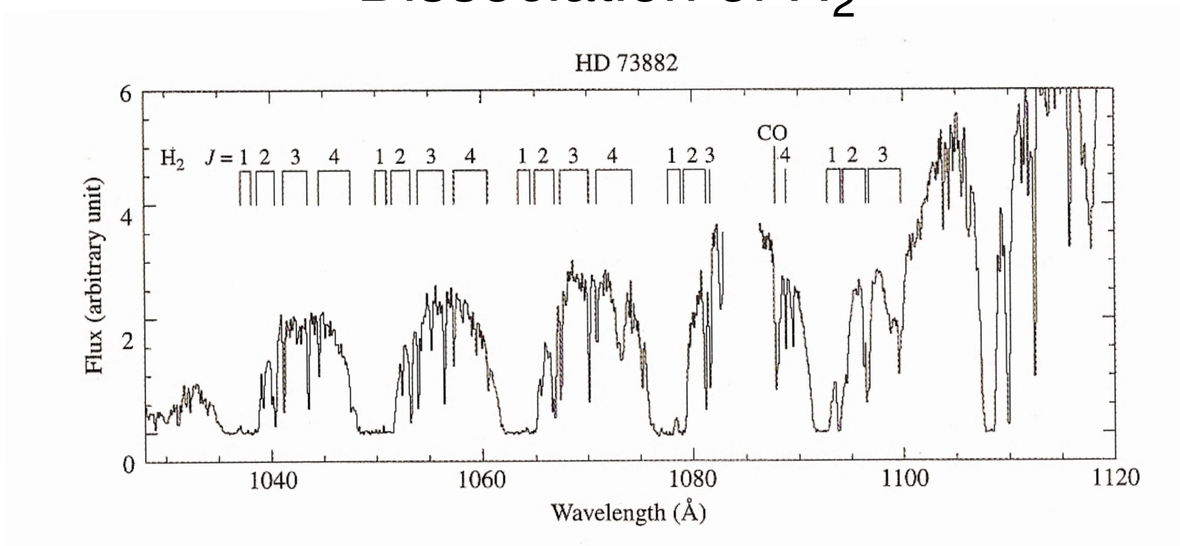
Figure from Cazuax & Spaans ApJ 611 40 2004

- Incident H atoms can be bound by long and short range forces, commonly called physi-adsortpion and chemi-adsorption.
- When another H strikes, it diffuses on the surface until it combines with a bound H, forms H_2, and leaves the surface.
- The processes of adsorption, sticking, binding, diffusion, formation and desorption depend on the gas and dust temperatures and on grain surface.
- Cazaux & Tielens (ApJ 575 L29 2000, 604 222 2004) find that the formation efficiency is ~ 1 for $T <$ 30K and ~ 0.3 for 100 $< T <$ 1000 K.

formation timescale: $\tau_{H_2} \approx 10^9 n_H^{-1}$ years

(low T; longer at higher T)

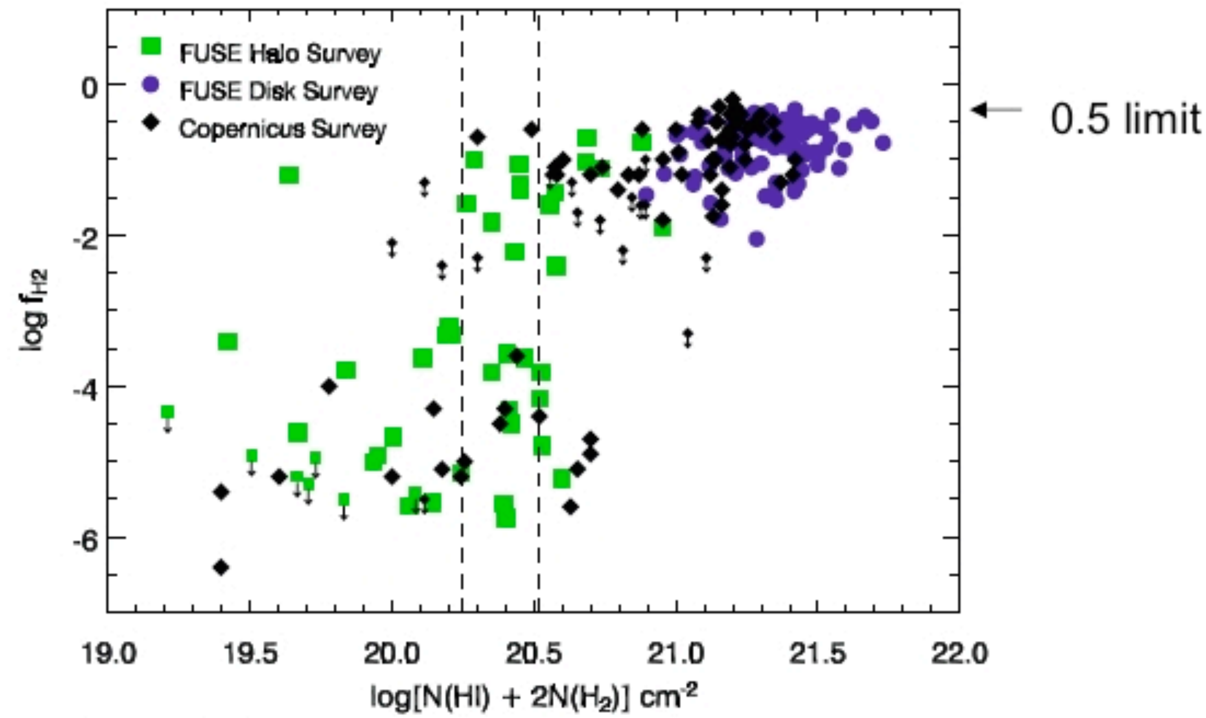
Dissociation of H_2



- collisional dissociation at T > 4000K
- direct photodissociation unlikely because it requires $h\nu > 14.7$ eV
- primarily by absorption of UV photons in the Lyman-Werner bands - excitation into positive energy states. Followed by radiative deexcitation. Most of the time, goes back to the ground state; about 13% of the time, dissociation results.
- H_2 optical depths are very large - “self-shielding”. Typically lines lie on damping part of curve of growth; thus rate $\propto N_H^{-1/2}$.

FUSE and Copernicus Survey Results

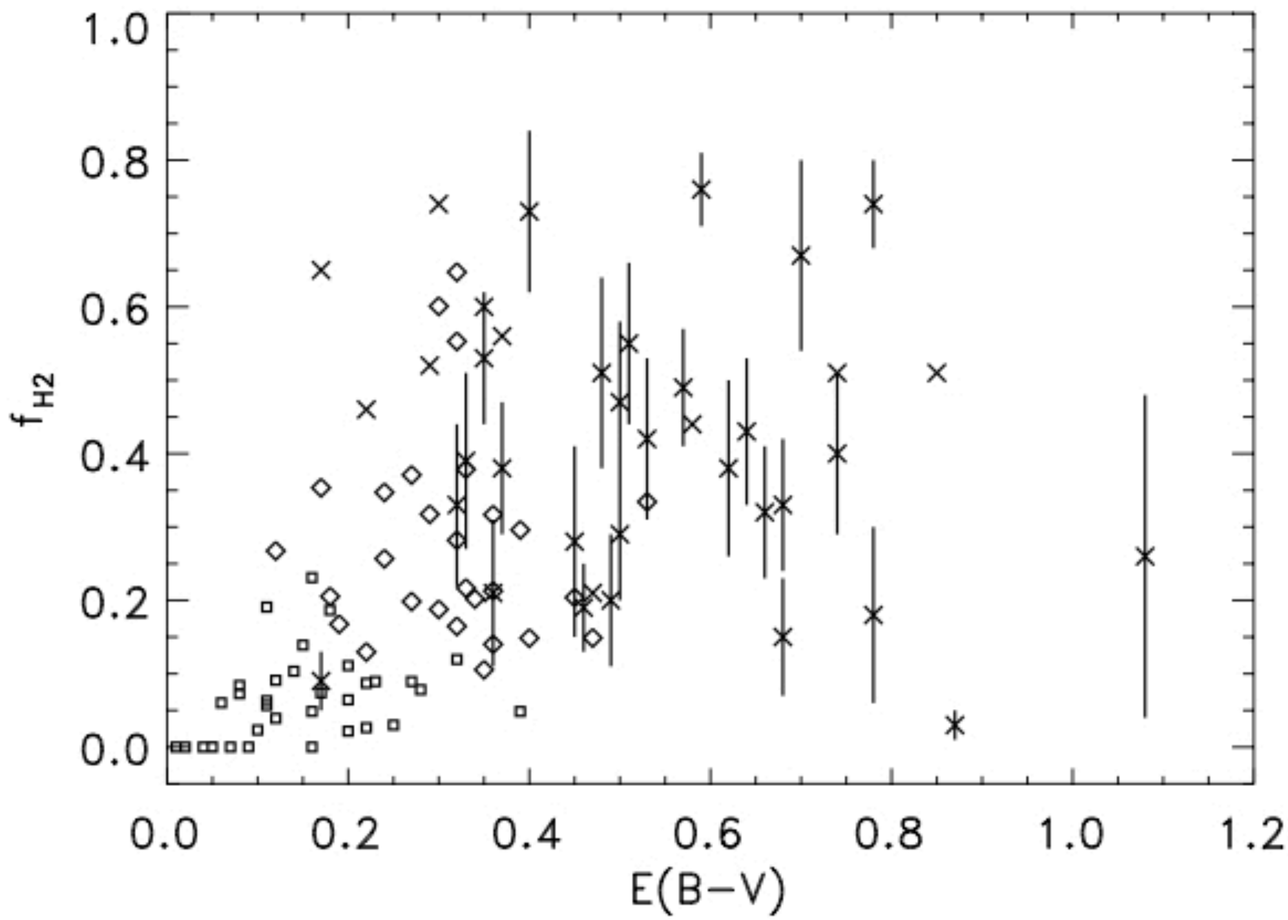
Gillmon et al (2006)



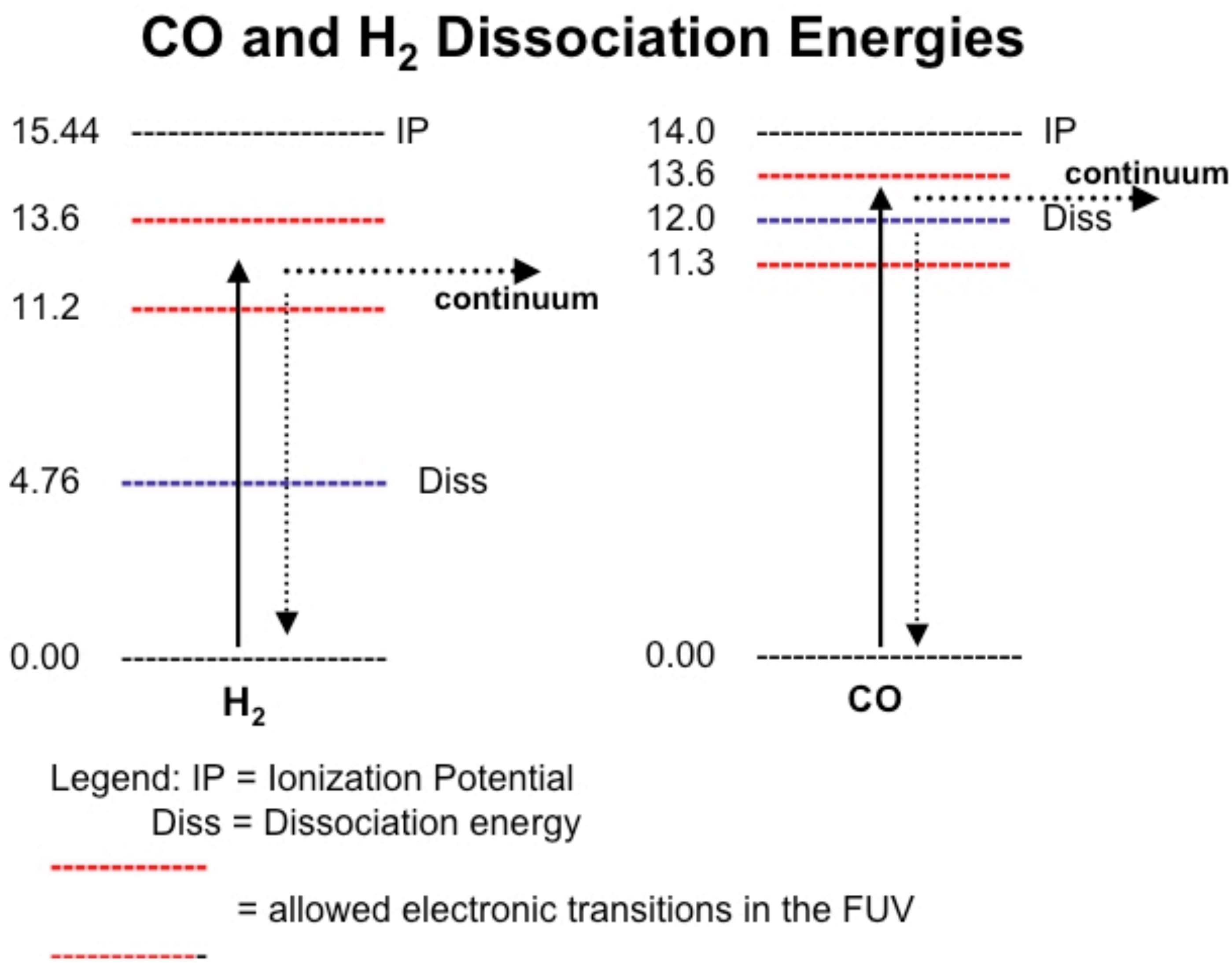
$$F = \frac{2N(\mathrm{H}_2)}{N(\mathrm{H}) + 2N(\mathrm{H}_2)} \quad \text{vs.} \quad N(\mathrm{H}) = N(\mathrm{H}) + 2N(\mathrm{H}_2)$$

Transition occurs near $N_H \sim 3 \times 10^{20}$ cm^{-2} :
$A_V \sim 0.15$ mag

Rachford et al. 2009



Transition occurs near : $A_V \sim$ 0.3-0.6 mag



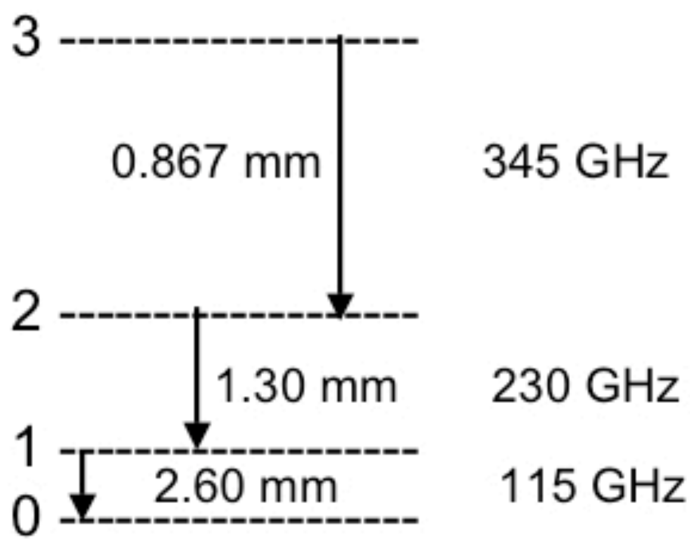
CO dissociates via FUV electronic transitions ala H_2; need shielding

Schematic "layers" of molecular cloud

UV →

H I; T >~ 30K; $N_H \sim 3 \times 10^{20}$ cm^{-2} ; $A_V \sim 0.15 – 0.3$	H_2; T <~ 20 K; $N_H \sim 3 \times 10^{20}$ cm^{-2} ; A_V >~ 0.3	CO; T <~ 15 K; $N_H \sim 3 \times 10^{20}$ cm^{-2} ; A_V >~ 0.5 - 1
photoelectric heating by UV on grains	UV attenuated by self-shielding	UV attenuated by H_2 , dust

CO rotational transitions; ideal for molecular cloud detection at mm wavelengths

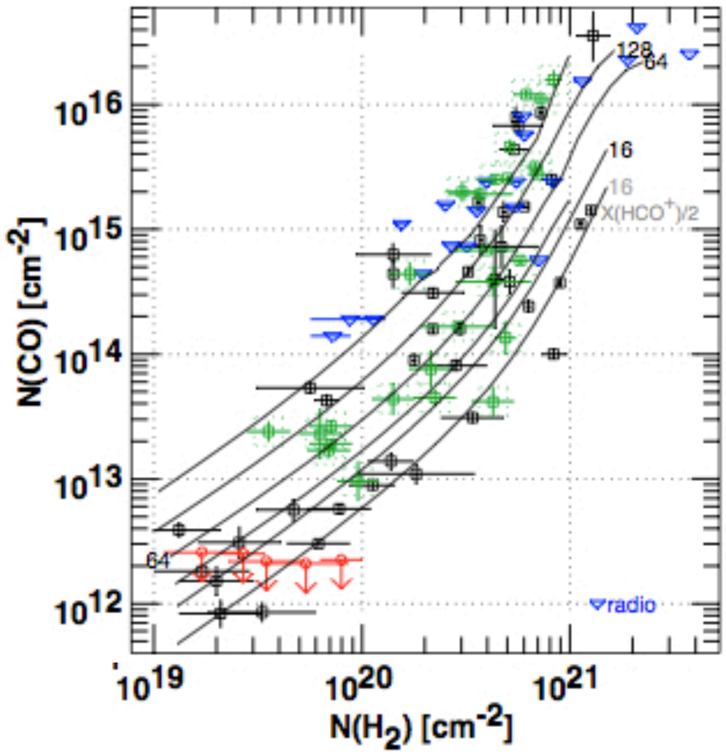


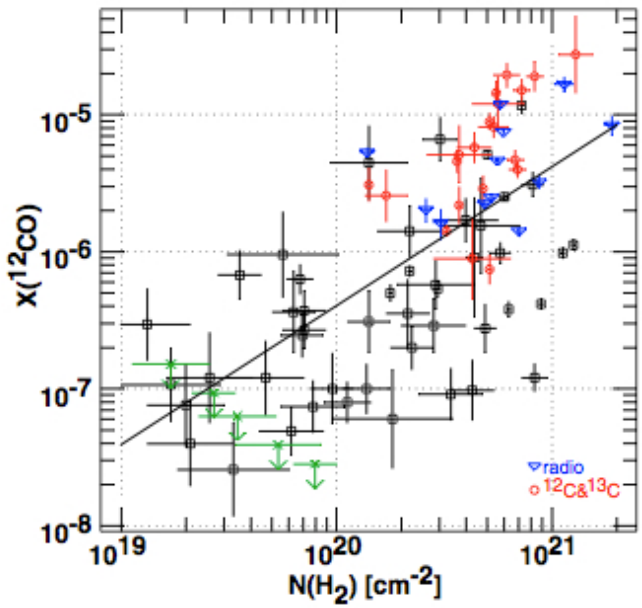
J = 1; E/k ~ 5.5 K
= 2; E/k ~ 16.5 K
= 3; E/k ~ 33 k

$A_{10} = 7.17 \times 10^{-8}\ s^{-1}$ $A_{J,J-1} = 3J^4/(2J+1)\ A_{10}$

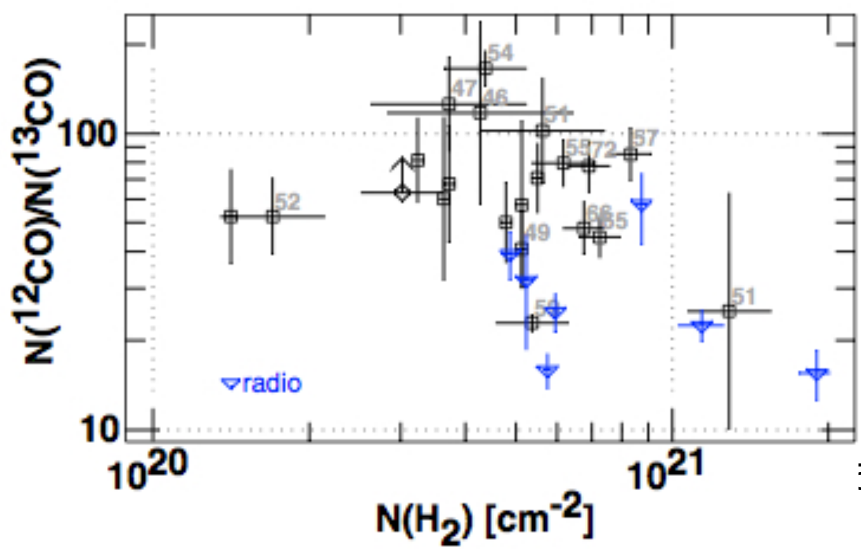
1-0, 2-1 main lines; 3-2 warmer regions
1-0 particularly likely to be optically thick.

Liszt 2007, A&A 476, 291 (translucent clouds)





CO becomes important tracer at $A_V \geq 1$



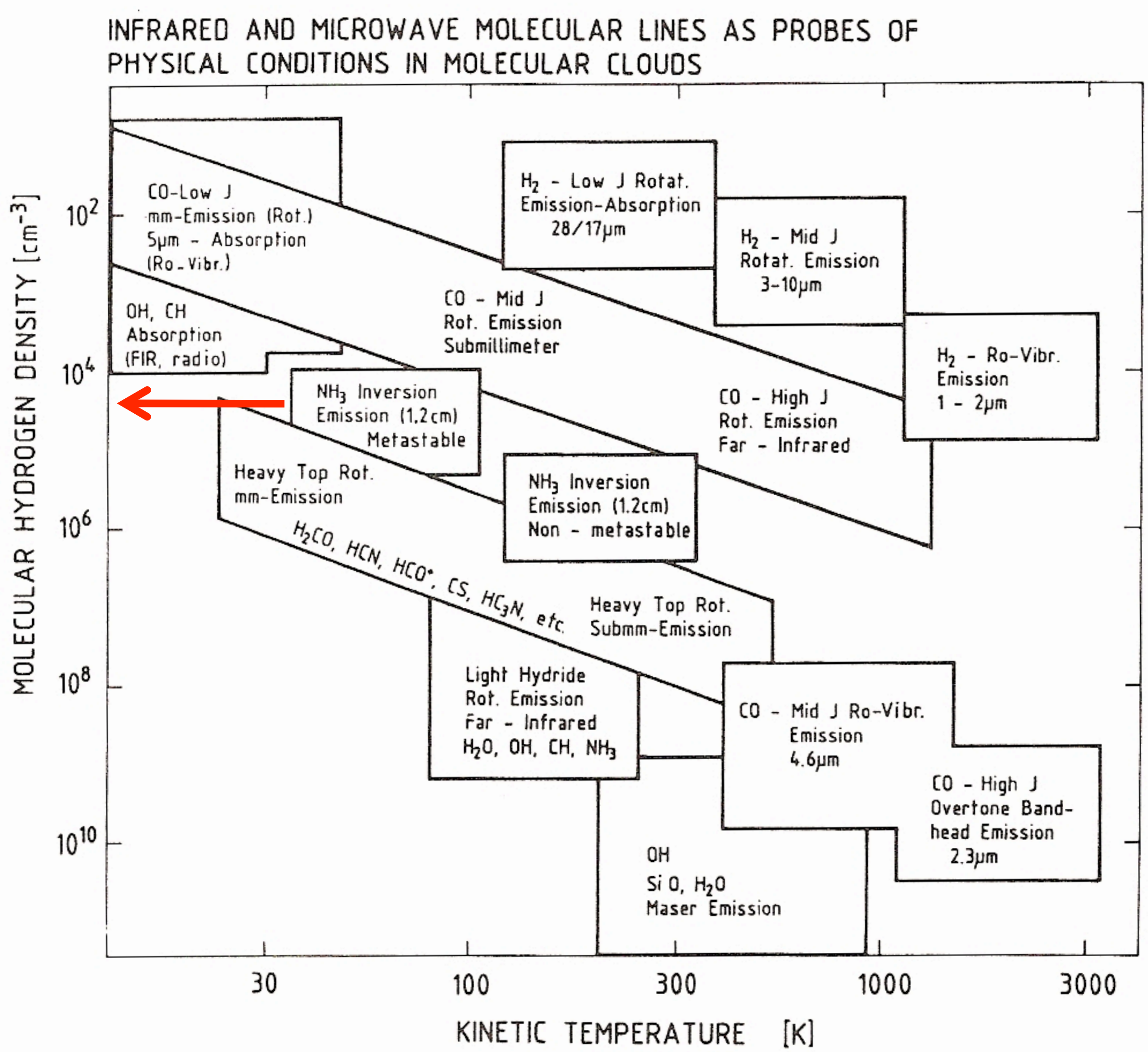


Figure 5. Molecular lines as probes of physical conditions in molecular clouds

Genzel 1990, NATO ASI 342 (Lada & Kylafis, eds)

Protostellar cloud cores and filaments

Cores do fragment in elongated structures along *infinite* filaments (see supplemental material); but they aren't infinite…

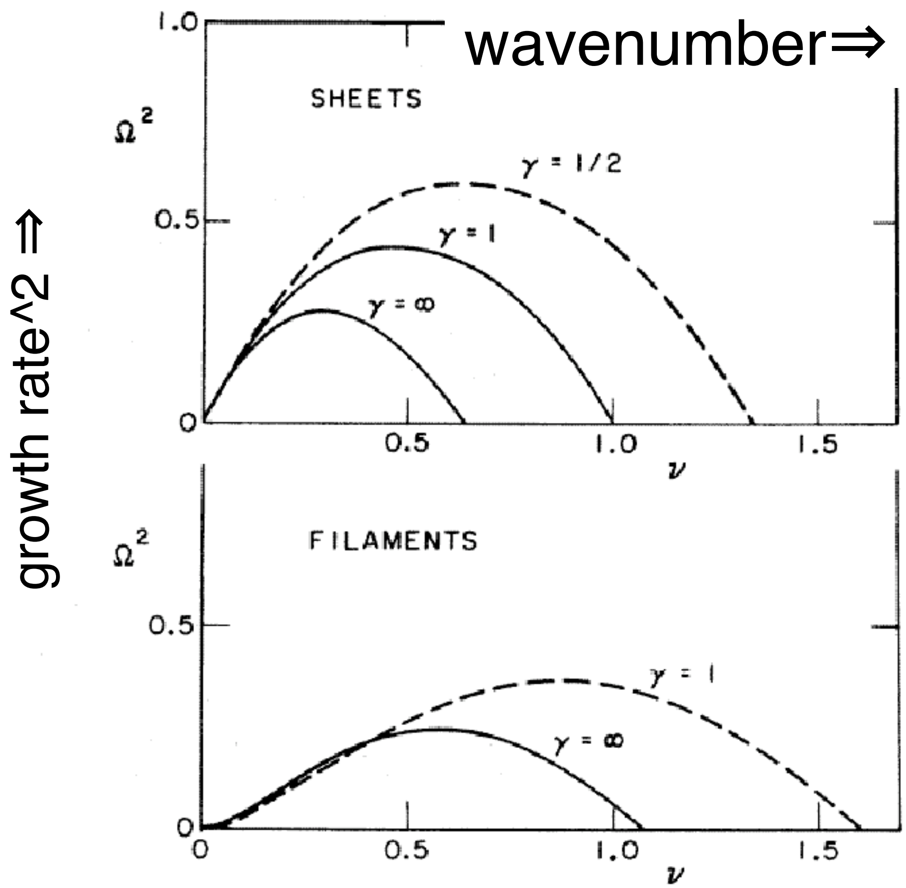


Figure 1. Dispersion relations for growing modes in polytropic sheets and filaments; the dimensionless squared growth rate Ω^2 is plotted versus the dimensionless wavenumber ν. For sheets, $\Omega = i\omega H/c(0)$ and $\nu = kH$, where H is defined by equation (14); for filaments, $\Omega = i\omega R/c(0)$ and $\nu = kR$, where R is defined by equation (20). The dashed curves are approximate dispersion relations estimated on the basis of scaling, as described in Sections 2.4 and 2.5.

One situation for which the assumption of a constant uniform background density as an equilibrium state is correct is the special case of a uniform, infinite, and infinitely-thin sheet of matter (or a thin sheet where gradients in the perpendicular (z) direction can be neglected). Then the conservation equations can be integrated vertically (in z), and the surface density Σ takes the place of the volume density;

$$\frac{\partial}{\partial t}\Sigma + \nabla \cdot (\Sigma \delta \mathbf{v}) = 0\,, \tag{A2.9}$$

$$\Sigma \frac{\partial}{\partial t}\delta \mathbf{v} = -\nabla \mathbf{P} - \mathbf{\Sigma} \nabla \mathbf{\Phi}\,, \tag{A2.10}$$

where the pressure has been vertically integrated. Linearizing as before, and starting from rest,

$$\frac{\partial}{\partial t}\delta\Sigma + \Sigma_{\circ} \nabla \cdot \delta \mathbf{v} = \mathbf{0}\,, \tag{A2.11}$$

$$\Sigma_{\circ} \frac{\partial}{\partial t}\delta \mathbf{v} = -c_s^2 \nabla \delta\Sigma - \Sigma_{\circ} \nabla \Phi\,. \tag{A2.12}$$

We assume here implicitly that the gradient in the background state pressure balances the gradient of the background potential. These gradients must be in the z direction by symmetry. We can use this fact with Gauss' law to derive the potential, by integrating the surface field $-\nabla\Phi$ over a pillbox of area $2dA$ encompassing the sheet at $z = 0$; then

$$2\,(-\nabla\Phi)\,dA = -4\pi G \Sigma\, dA\,, \tag{A2.13}$$

so that

$$\nabla \Phi = 2\pi G \Sigma\,. \tag{A2.14}$$

Therefore,

$$\lim_{\epsilon \to \infty} \int_{-\epsilon}^{+\epsilon} dz\, \nabla^2 \Phi \;=\; \lim_{\epsilon \to \infty} \left[\nabla\Phi \,|_{+\epsilon} \;-\; \nabla\Phi \,|_{-\epsilon} \right] \;=\; 4\pi G \Sigma \,. \tag{A2.15}$$

From basic definitions, this implies

$$\nabla^2 \Phi \;=\; 4\pi G \Sigma \delta(z) \,, \tag{A2.16}$$

where δ is the Dirac delta function.

As before, we analyze behavior for a plane wave perturbation in the x direction,

$$\delta\Phi \;=\; \delta\Phi_a \, e^{i(kx \,-\, \omega t)} \,, \tag{A2.17}$$

where $\delta\Phi_a$ is a (constant) amplitude. This equation can only apply in the plane $z = 0$; for $z \neq 0$,

$$\nabla^2 \delta\Phi \;=\; 0 \,. \tag{A2.18}$$

There is only one continuous function which satisfies both these constraints, and also goes to zero at large z,

$$\delta\Phi \;=\; \delta\Phi_a \, e^{i(kx \,-\, \omega t) - |kz|} \,. \tag{A2.19}$$

Returning to the Gauss' law evaluation,

$$\lim_{\epsilon \to \infty} \left[\frac{\partial}{\partial z} \delta\Phi \,|_{+\epsilon} \;-\; \frac{\partial}{\partial z} \delta\Phi \,|_{-\epsilon} \right] \;=\; -2 \,|\, k \,|\, \delta\Phi_a \;=\; 4\pi G \delta\Sigma \,, \tag{A2.20}$$

Returning to the Gauss' law evaluation,

$$\lim_{\epsilon \to \infty} \left[\frac{\partial}{\partial z} \delta\Phi \mid_{+\epsilon} \; - \; \frac{\partial}{\partial z} \delta\Phi \mid_{-\epsilon} \right] \; = \; -2 \mid k \mid \delta\Phi_a \; = \; 4\pi G \delta\Sigma \, , \qquad (A2.20)$$

so that

$$\Phi_a \; = \; -2\pi G \delta\Sigma / \mid k \mid \, . \qquad (A2.21)$$

Substitution of this result into the linearized equations leads to the dispersion relation

$$\omega^2 \; = \; c_s^2 k^2 \; - \; 2\pi G \Sigma_\circ \mid k \mid \, . \qquad (A2.22)$$

Again, when $\omega^2 < 0$, exponential growth occurs, so the sheet is gravitationally unstable for

$$\mid k_J \mid < 2\pi G \Sigma_\circ / c_s^2 \, . \qquad (A2.23)$$

Thus, just as in the basic Jeans mass calculation, wavelengths greater than $\lambda_J > 2\pi/k_J$ are unstable. However, the sheet dispersion relation differs in an important way from the basic Jeans mass result. Defining a growth rate $\Gamma = -i\omega$, and taking only real positive k,

$$\Gamma^2 \; = \; 2\pi G \Sigma_\circ k \; - \; c_s^2 k^2 \, , \qquad (A2.24)$$

so that $\Gamma \to 0$ as $k \to 0$. The growth rate has a maximum at $k_{max} = k_J/2$, unlike the basic Jeans mass calculation, where the growth rate is largest for $k \to 0$.

However; can't have infinite filaments and don't have uniform density spheres.

Consider a sphere of radius R with uniform density ρ. At large scales, where gas pressure is not important, all radii fall in at the same time. One can see this from dimensional analysis:

$$v^2 \sim GM/R \sim G\rho R^2\,; \qquad thus \qquad t_{ff} \sim R/v \sim (G\rho)^{1/2}\,. \tag{1}$$

However, the global nature of gravity in a non-uniform, non-spherical volume results in non-linear accelerations as a function of position, so that it is essentially impossible to prevent collapse somewhere (unless gravity is too weak or expansion is too large). For example, consider a thin uniform density circular sheet; the gravitational potential is

$$\Phi = -4G\Sigma R\, E(r/R)\,, \tag{2}$$

where Σ is the surface density and E is the second complete elliptic integral. The gravitational acceleration toward the center at r is

$$a_r = -\frac{\partial \Phi}{\partial r} = 4G\Sigma \frac{R}{r}\left[K(r/R) - E(r/R)\right]\,, \tag{3}$$

where K is the first complete elliptic integral. Expanding,

$$a_r = \frac{1}{2}\frac{dv^2}{dr} = \pi G\Sigma\left[\frac{r}{R} + \frac{3}{8}\left(\frac{r}{R}\right)^3 + \frac{45}{192}\left(\frac{r}{R}\right)^5 + ...\right]. \tag{4}$$

The figure shows this function, along with a comparable acceleration for a thin cylinder (filament).